

CHAPTER 1

SEEING THE SUN

1.1 Introduction

As seen from the Earth, the stars seem cold and remote. Yet one star is so blindingly present in our lives that many people do not realize that it is a star at all. That star is the Sun, and it is the subject of the first two chapters of this book. This chapter concentrates on the observed properties of the Sun, and as you will see, the majority of observations depend in some way on the light or, more generally, the electromagnetic radiation that the Sun emits. Hence a secondary theme of this chapter is to introduce some concepts about the nature of light that are vital not only for solar science, but for astronomy in general.

Given that the Sun is, by far, the closest star to the Earth, it offers us unique opportunities to carry out very detailed studies of its behaviour. We will review some of the most important features of the Sun by concentrating on those parts of the Sun that can be observed directly. This sets the scene for Chapter 2 in which we explore the physical mechanisms that give rise to the Sun's observed characteristics.

1.2 Seeing the Sun's surface

1.2.1 Introducing the photosphere

Figure 1.1 is a photograph of the Sun. At first sight it looks like a pretty ordinary astronomical photograph – you might even mistake the Sun's image for that of some other body, a planet perhaps, or even a moon. But the Sun is neither a planet nor a moon; it is a star – our star – the only star that is sufficiently close for us to be able to examine its visible surface in great detail. For this reason a full discussion of Figure 1.1 is a good starting point for a book dealing with the stars.

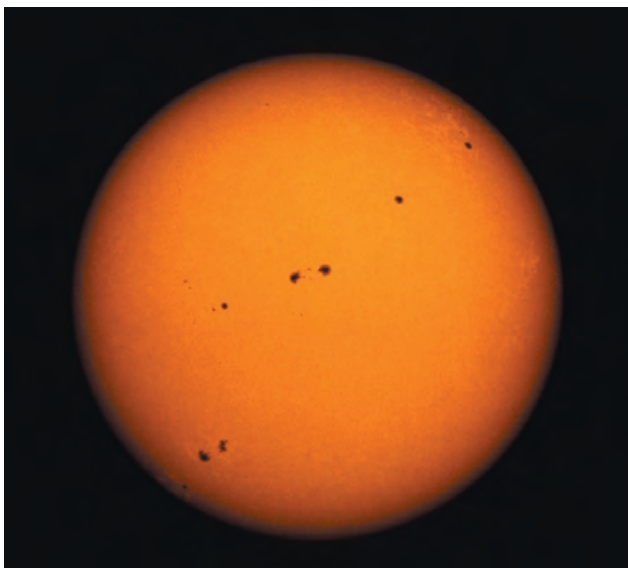


Figure 1.1 A photograph of the Sun. (NOAO)

There are a number of ways in which the Sun of Figure 1.1 differs from a planet or a moon. The most important of these differences is simply that the Sun is very much brighter than any planet or moon, from which it can be concluded that the Sun is a prodigious source of energy. The energy emitted by the Sun is crucial to our very existence, since it provides us with warmth and light. Without it, life might never have arisen on the Earth, and life as we know it today could certainly not be sustained. The rate at which energy is radiated by the Sun (i.e. the total amount of energy radiated per second) is called the **solar luminosity**. It is denoted by the symbol L_{\odot} , and is about $3.84 \times 10^{26} \text{ J s}^{-1}$. (The \odot symbol is an ancient astrological sign for the Sun. Despite the distaste with which most astronomers view astrology, the symbol is widely used to denote ‘solar’ quantities, i.e. quantities pertaining to the Sun.)

QUESTION 1.1

Note that a table of often-used values is given in Appendix A1.

What is the value of the solar luminosity, L_{\odot} , in terms of the SI unit of power – the watt (W)? Given that a typical large power station produces energy at the rate of $2.5 \times 10^9 \text{ W}$, work out the number of such power stations that would be required to match the energy output of the Sun.

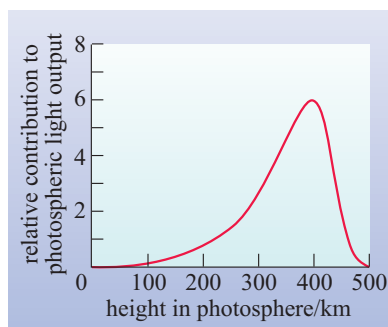


Figure 1.2 The relative amounts of photospheric light originating at various heights in the photosphere. Heights are measured from a precisely defined reference level that roughly corresponds to the greatest depth that can be ‘seen’.

A second major difference between the Sun and the planets that orbit it concerns size. The visible disc portrayed in Figure 1.1 is actually about $1.4 \times 10^6 \text{ km}$ in diameter. That’s nearly ten times greater than the diameter of Jupiter, the largest planet, and about a hundred times greater than the diameter of the Earth. Clearly, the Sun is very much larger than any planet in the Solar System.

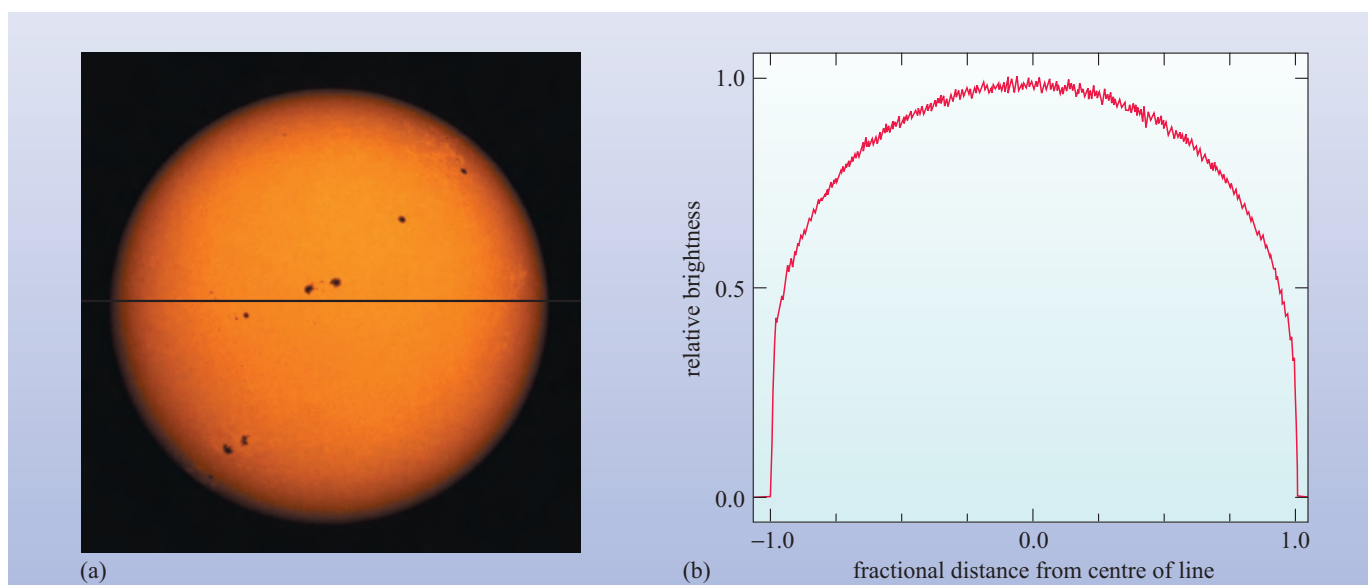
A third difference between a picture of the Sun and a picture of a body like the Earth or the Moon is that the visible surface of the Sun is not really a surface at all, but rather a thin, semi-transparent shell of gaseous material. Photographing the Sun is rather like photographing a cloud or a bank of fog; the light that makes up the photographic image comes from a range of depths and not from a single well-defined surface. In the case of Figure 1.1 almost all the light comes from a layer about 500 km thick called the **photosphere** (meaning ‘the sphere of light’). Now, 500 km may sound pretty thick but, remember, the Sun is about 1.4 million kilometres across, so in comparison with the total solar diameter the photosphere really is a thin spherical shell. When we photograph the Sun we see into the photosphere but not through it. Figure 1.2 gives a rough idea of the relative proportions of photospheric light coming from different parts of the photosphere.

In visual terms, the photosphere is the closest thing to a surface that the Sun has to offer, but in physical terms the photosphere is much more like an atmospheric layer. In fact, the air you are breathing right now is more than a thousand times denser than the material that makes up the photosphere. If a space probe were sent into the photosphere the frictional resistance it would encounter would be almost negligible. A far greater impediment to such a mission would be the temperature. The photosphere gives off light because it is hot. Typical temperatures range from about 9000 K in the lowest parts of the photosphere to about 4500 K at the top. Most of the photospheric light comes from a region where the temperature is between 5800 K and 6000 K, so it is conventional to use values such as these to represent the ‘surface temperature’ of the Sun: however, the term is not really very meaningful and must be treated with some caution.

1.2.2 Large-scale features of the photosphere

Now that you are acquainted with the broad features of the photosphere (size, thickness, temperature, gaseousness) it makes sense to take a more detailed look at Figure 1.1. There are two main points to note. First, it should be fairly obvious that the edges of the photosphere are darker than the centre: there is a steady reduction in brightness as the distance from the centre of the image increases. Because the edge of the solar image is known as the **solar limb**, this gradual fall-off in brightness is called **limb darkening**. The second feature to note is that the photosphere is marked by dark blotches called **sunspots**. It turns out that both these features are important in understanding the Sun, so we shall discuss them in turn, starting with limb darkening.

Although limb darkening can be seen in Figure 1.1, it is more clearly demonstrated by using a graph to show the relative brightness of points along a line that crosses the solar disc, passing through the centre. Such a line is shown in Figure 1.3a and the corresponding graph in Figure 1.3b. As you can see from the graph, the brightness falls off very rapidly near the limb. This means that the solar limb is quite well defined, which is important because it makes it possible to use the location of the limb in various kinds of measurement.



Now, why should the limb of the Sun be darkened in the way indicated by Figure 1.3? In order to understand this, it is again necessary to recognize that the solar photosphere is something we look into rather than something we simply look at. As a first step towards understanding limb darkening, let us consider a scenario that is less complex than the Sun. In this simplified case we have a spherical cloud of gas in which the gas has uniform density. When we look at this cloud, our line of sight penetrates a certain distance into the cloud. Because this cloud has a uniform density, the distance along the line of sight from the edge of the cloud to the furthest point that we can see is the same for any line of sight. However, as Figure 1.4a shows, when we look along different lines of sight into the cloud, the depth below the surface of the furthest point that we can see (i.e. as measured along a line from the surface of the cloud to its centre) is *not* constant.

Figure 1.3 (a) The visible solar disc, crossed by a straight line. (b) The relative brightness of the photosphere at various points along the straight line shown in (a). ((a) NOAO; (b) Foukal, 1990)

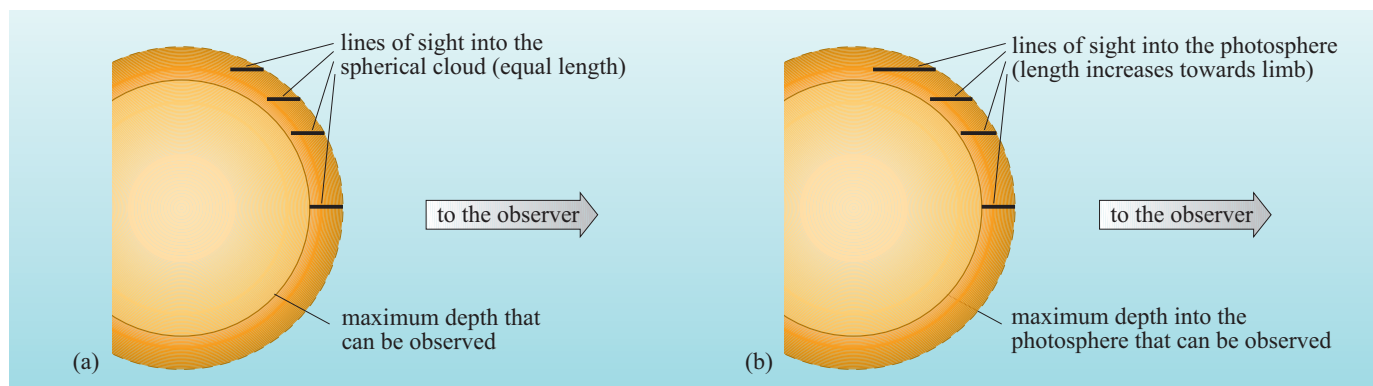


Figure 1.4 (a) For a gaseous spherical cloud of uniform density, we can see a uniform distance along the line of sight into the cloud. The depth below the surface of the furthest point that is visible becomes shallower as lines of sight approach the limb of the cloud. (b) In the Sun, the density changes with depth, and the lengths of lines of sight within the cloud increase as they approach the solar limb. However, the overall result is similar to case (a) in that the depth below the surface of the furthest point that we can see again becomes shallower as lines of sight approach the limb of the Sun. (Note that the relative distances that the lines of sight penetrate into the Sun are grossly exaggerated in this diagram.)

- From Figure 1.4a, at which position do we see to the greatest depth below the surface of the cloud, and at which position do we see to the shallowest depth below the surface of the cloud?
- The line of sight that penetrates to the greatest depth is that which is directed towards the centre of the cloud. The lines of sight that reach the shallowest depths are those near the limb of the cloud

So there is an important geometrical effect here; lines of sight near the centre of the disc penetrate to a greater depth than lines of sight that are nearer the limb of the disc. Let us now consider a situation in which the temperature of the gas drops with distance from the centre of the cloud. Lines of sight toward the centre of the cloud will penetrate deepest into the cloud and hence allow us to see the hottest regions that are visible. Towards the limb, the lines of sight will only reach shallower and hence cooler regions of the cloud. An image of this cloud would then show limb darkening because the cooler material seen close to the limb simply gives off less visible light than the hotter material seen at the centre.

The situation in the Sun is somewhat more complex than this simplified scenario. While the temperature of the photosphere of the Sun does indeed increase with depth, the density of gas also increases with depth. This change in density has the effect that the distance along a line of sight from the surface to the furthest point that we could see is not constant. In fact, this distance actually increases as lines of sight move away from the centre of the solar disc. However, it turns out that despite this increase in distance of the line of sight within the photosphere, the geometrical effect that you have already seen in the simplified case is more important. This situation is illustrated schematically in Figure 1.4b. The overall result is that lines of sight towards the centre of the solar disc penetrate to a depth of about 500 km below the top of the photosphere, while lines of sight towards the solar limb penetrate to shallower depths. Because, as has already been pointed out, the temperature of the photosphere increases with depth, this variation in the depth to which we can see gives rise to limb darkening across the solar disc.

Turning now to the sunspots seen in Figure 1.1, it's worth noting straightaway that their darkness is also a consequence of temperature. Sunspots are large, relatively cool regions of the photosphere; the temperature at the centre of a sunspot is typically 4200 K, which is much less than the 6000 K or so of the surrounding photosphere. Consequently, sunspots are seen as dark patches against the bright background of the photosphere.

QUESTION 1.2

Using Figure 1.1, roughly estimate the diameter of a large sunspot.

Warning: do not attempt to look directly at the Sun.

Sunspots have been observed since ancient times, but their serious study really began in 1610, or shortly thereafter, when telescopes were just beginning to be used for astronomical purposes. Early solar observers, such as Galileo Galilei (1564–1642), David Fabricius (1564–1617) and the appropriately named Christoph Scheiner (1575–1650), soon discovered that sunspots appeared to move across the face of the Sun. This was eventually accepted as clear evidence that the Sun rotates, carrying the sunspots with it as it turns on its axis.

Individual sunspots are transient phenomena, but their lives are sufficiently long – typically a few weeks – that it is often possible to observe them crossing the entire solar disc. Sometimes, particularly long-lived spots, or groups of spots, can even be seen re-appearing over the limb of the Sun after they have crossed the far side

GALILEO GALILEI (1564–1642)

Galileo Galilei (Figure 1.5), who is usually referred to solely by his first name, was born in Pisa in 1564. He studied medicine but did not take his degree. Instead he developed an interest in mathematics and physics which he studied at home. In 1589 he was appointed professor of mathematics at Pisa, and three years later moved to a similar post in Padua. Galileo had many interests and arguably his most significant work was that which laid the foundations for the scientific study of motion.

Although Galileo did not invent the telescope, he was the first person to recognize its potential in astronomy. After learning of the telescope in 1609 he soon built one of his own and began making observations of the Sun, planets and stars. He reported his discoveries in a book published in 1610 called *Siderius Nuncius* (*The Starry Messenger*) which brought him widespread fame and the opportunity to advance his career under the patronage of the Grand Duke of Tuscany in Florence. However, controversy soon followed, when in 1613 he publicly supported the Sun-centred model of the Solar System that had been put forward by Nicolaus Copernicus (1473–1543). The Church deemed this view heretical, and Galileo then became embroiled in a long-running dispute over his beliefs. Events came to a head in 1633, when, under the threat of torture from the inquisition, Galileo was forced to reject the Copernican view, and was kept under house arrest for the remainder of his life. Following a campaign by prominent astronomers and historians of science, the sentence that had been passed on Galileo was formally retracted in 1992.

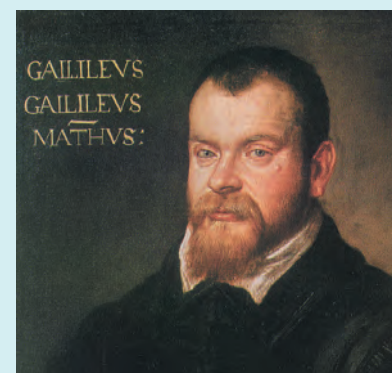


Figure 1.5 Galileo Galilei.

of the Sun. A sequence of photographs illustrating this effect is shown in Figure 1.6. In this particular case the spots were observed to take about 28 days to make a complete circuit. You might well think that this implies a 28 day rotational period for the Sun, or at least for the photosphere, but things are not quite so simple. In the first place, because the Earth orbits the Sun once a year (moving around the Sun in the same sense that the Sun rotates around its own axis), the rotational period observed from Earth is actually slightly longer than the **sidereal period**; that is, the period as measured with respect to the stars.



Figure 1.6 A sequence of images showing the apparent motion of sunspots across the face of the Sun. Note that each image is dated and that the whole sequence covers a period of about five weeks. (SOHO (ESA and NASA))

Second, owing to the fact that the Sun is a gaseous body, the nature of its rotation is very different from the rotation of a solid body such as the Earth. On a solid body everything ‘rotates together’, each part of the surface keeping in step with every other part, but on a gaseous object it is quite possible for different parts to rotate at different angular speeds, and this is just what happens on the Sun. Studies of sunspots and other indicators of **solar rotation** show that points on the solar equator have a sidereal period that is just under 26 days, whereas points further north or south have considerably longer periods: more than 26 days at a latitude of 30° , and about 30 days at a latitude of 60° . It is actually very difficult to measure the rotational period close to the poles, but it seems to be about 36 days. This rather complicated state of affairs is described by saying that the Sun exhibits **differential rotation**. More precise information about the varying rate of photospheric rotation is given in Figure 1.7.

Apart from their role as tracers of rotation, sunspots are also good indicators of another large-scale phenomenon: **solar activity**. Data collected over many decades (see Figure 1.8) clearly show that the fraction of the solar disc covered by sunspots changes with time in a more or less regular way. A period of roughly 11 years separates each occurrence of maximum coverage, and hence of maximum solar activity, from its successor. Images of the Sun recorded at a time of maximum activity and at a time when activity is minimal differ markedly – as can be seen from Figure 1.9. Many other solar phenomena, some of which will be discussed later, also participate in the 11-year **solar activity cycle**, but none is as easy to observe as sunspots.

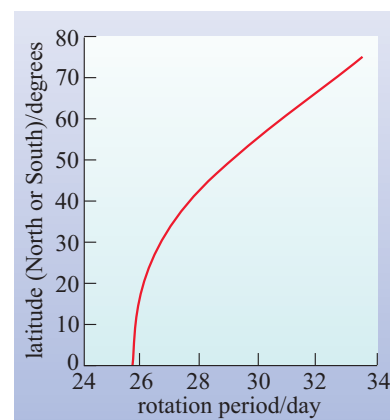


Figure 1.7 The sidereal period of the solar photosphere at various latitudes.

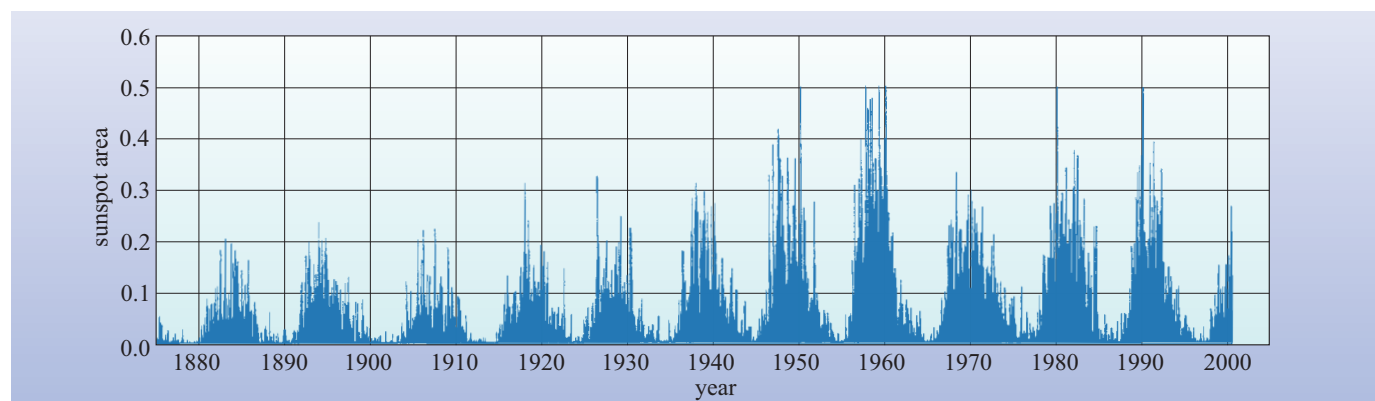


Figure 1.8 The 11-year variation of solar activity with time, as indicated by the percentage of the area of the Sun’s visible disc covered by sunspots. (D. Hathaway, NASA/MSFC)

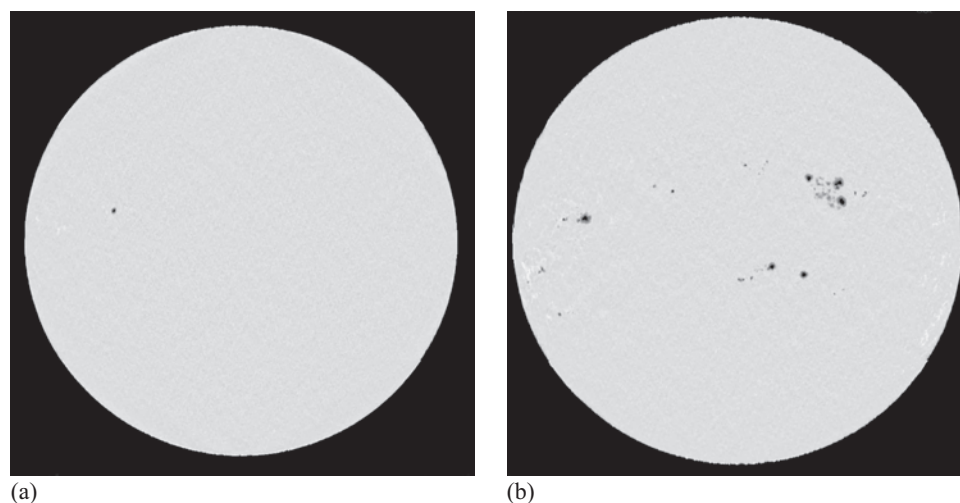


Figure 1.9 The solar photosphere, at times of low and high solar activity. (a) Low solar activity (22 June 1996). (b) High solar activity (23 December 2001). (Note that these images have been processed to remove the effects of limb darkening.) (Data provided by NSO)

1.2.3 Small-scale features of the photosphere

Although the origin of the solar activity cycle remains one of the Sun's outstanding mysteries, a good deal is known about sunspots. For example, it is well established that sunspots have stronger magnetic fields and different patterns of motion from their surroundings. Knowledge of this kind partly results from the study of magnified views of localized regions of the photosphere rather than photographs of the full solar disc. A magnified view of just this kind, showing the detailed structure of a sunspot, is shown in Figure 1.10. Such views are of great importance; the Sun is the only star sufficiently close to allow such detailed imaging of its surface.

Another small-scale phenomenon that is thought to be common in stars, but which can actually be seen only in the Sun's photosphere, is shown in Figure 1.11. The figure provides an instantaneous snapshot of the **solar granulation** – a seething pattern of bright cell-like **granules** that covers the photosphere. Each granule is typically about 1000 km across and lives for five to ten minutes. Detailed studies of the granules show that they are the tops of rising columns of hot material coming from deeper regions of the Sun where the temperatures are higher. The rising material travels upwards at a speed of 1 km s^{-1} , or thereabouts, and then spreads out horizontally, radiating away its excess thermal energy. The dark 'lanes' between granules are regions where the cooled material descends back into the solar interior. The significance of these motions will be more fully explored in Section 2.2.

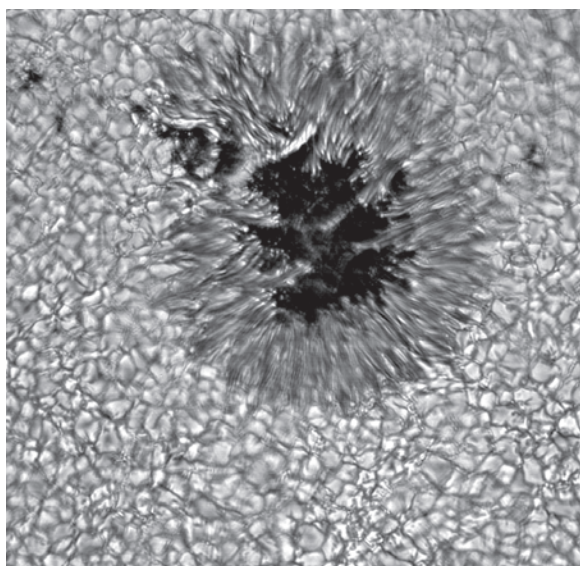


Figure 1.10 A highly magnified view of a sunspot. The sunspot itself consists of the very dark area and the pattern of fronds that seem to emanate from this region. The 'background' pattern on which the sunspot is located is the solar granulation. This image shows a region that is approximately $75\,000 \text{ km} \times 75\,000 \text{ km}$ in extent. (G. Scharmer, Royal Swedish Academy of Sciences)

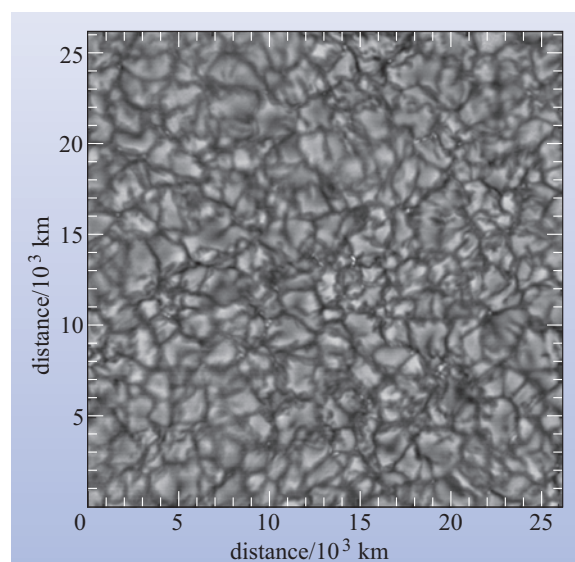
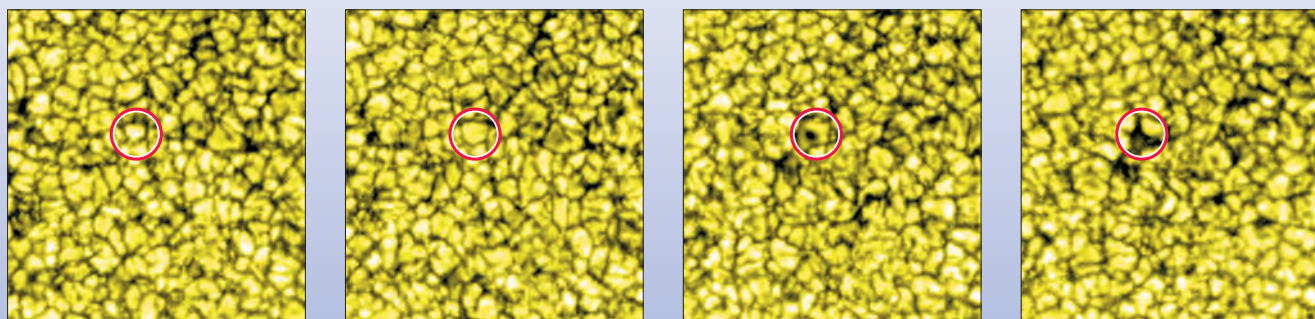


Figure 1.11 A highly magnified view of a small part of the photosphere clearly showing the solar granulation. (NOAO)

The blurring effects of the Earth's atmosphere make it difficult to carry out detailed studies of the solar granulation. Nonetheless, it is sometimes possible to obtain sequences of photographs that clearly show the formation or disappearance of individual granules. Such a sequence is shown in Figure 1.12.



1.2.4 The nature of light

Note: throughout this chapter the term 'light' refers specifically to 'ordinary' visible light.

Although our discussion of the photosphere has concentrated on its visual appearance, it is important not to forget that the photosphere is the source of the light that illuminates the Earth and, through the effect of photosynthesis in plants, is the source of energy that keeps us alive. From our human perspective, light is the main product of the photosphere and will be a major concern throughout this book, so this is a good point at which to gather together a number of basic facts about the nature of light. Because these facts constitute 'essential scientific background' rather than a continuation of the astronomical storyline we have been developing so far, they will be separated from the rest of the section by enclosing them in a blue toned box. Such boxes will be used throughout this book to enable you to identify items of background science wherever they arise. One such example is Box 1.1.

Figure 1.12 A sequence of photographs of the solar granulation. Note that one particular granule has been highlighted by a red circle. This granule grows and decays over the 8 minute interval covered by the photographs. (BASS 2000)

BOX 1.1 THE NATURE OF LIGHT

The electromagnetic wave model of light

It is well known that a magnet is able to influence certain objects (e.g. other magnets) without touching them. This phenomenon is 'explained' by saying that the magnet produces a **magnetic field**, which occupies the space around the magnet and gives rise to the forces that act on the affected objects. To account for these forces, the magnetic field at any point must have both a strength and a direction. Consequently, the magnetic field at any point can be represented by an arrow, since an arrow has a length that can represent the strength of the field, and an orientation that can be made to correspond to the direction of the field. The use of arrows to represent a magnetic field is illustrated in Figure 1.13a. In a similar way, a suitable distribution of

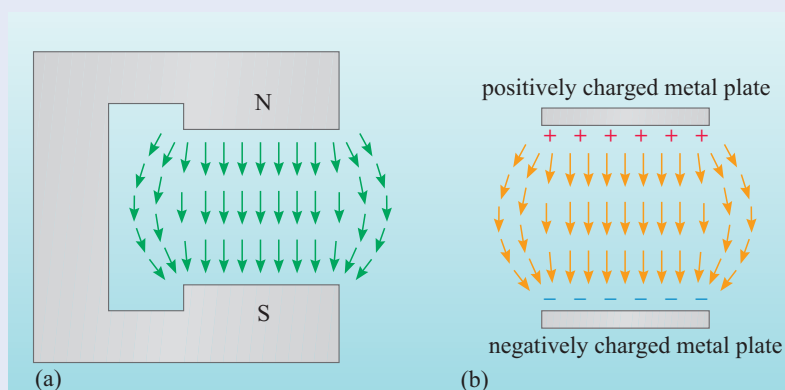


Figure 1.13 (a) A magnetic field. (b) An electric field.

positive (+) and negative (–) electric charges will give rise to an **electric field**, which can also be represented, at any point, by an arrow of appropriate length and orientation, as in Figure 1.13b.

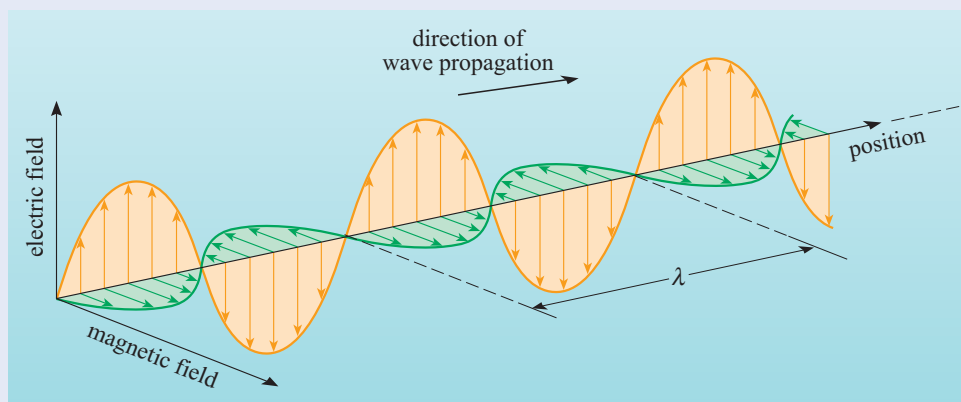


Figure 1.14
An electromagnetic wave.

In the 1860s, while carrying out a mathematical investigation of electricity and magnetism, the Scottish physicist James Clerk Maxwell showed that it was possible to create self-sustaining patterns of fluctuating electric and magnetic fields that could oscillate together through space at a certain pre-determined speed. These fluctuating field patterns are called **electromagnetic waves**, and an instantaneous ‘snapshot’ of a small part of the simplest such wave is shown in Figure 1.14. The electric and magnetic fields that make up the wave are always at right angles to each other, and both fields are at right angles to the direction in which the wave is travelling. Remember, Figure 1.14 is only a snapshot, so you should imagine the whole pattern moving in the direction of travel rather like an ocean wave moving across open water. Maxwell found that the speed at which electromagnetic waves had to travel was very close to the best values then available for the speed of light, so he suggested that rays of light were nothing other than electromagnetic waves.

An important characteristic of any electromagnetic wave is the distance between successive maxima of the electric or magnetic field (that is, the distance from one peak of the wave to the next). This distance is called the **wavelength** and is denoted by the Greek letter λ (pronounced ‘lambda’). The wavelengths of all forms of visible light are very tiny, and different wavelengths correspond to different colours. Red light has a wavelength of about $700 \times 10^{-9} \text{ m}$ ($= 700$ nanometres, nm), and the wavelength of violet light is around $420 \times 10^{-9} \text{ m}$ ($= 420$ nm). White light, which is a mixture of all the possible colours, contains all the wavelengths between these rough limits. (You will find more on this subject in Section 1.3.2.)

Imagine yourself observing a wave like that illustrated in Figure 1.14 as it moves past some fixed point. If the

speed of the wave is v and its wavelength is λ , then you should be able to convince yourself that the number of wavelengths that will pass the fixed point in a second is just v/λ . This quantity, the number of wavelengths passing a fixed point in one second, is called the **frequency** of the wave; it is measured in SI units called **hertz (Hz)** (equivalent to s^{-1}) and is denoted by the letter f . Thus, for any wave, $f = v/\lambda$. Or, more conventionally,

$$v = f\lambda \quad (1.1)$$

For electromagnetic waves the value of v depends on the **medium** (e.g. air, glass and water) through which they move. The maximum value of v occurs when the waves travel through a vacuum (that is, empty space). Under these conditions the speed of the waves is very nearly $3.00 \times 10^8 \text{ m s}^{-1}$ (the exact value is $2.997\,924\,58 \times 10^8 \text{ m s}^{-1}$). This quantity is of such importance that it is given its own symbol, c , and is referred to as the **speed of light in a vacuum**. Thus, for electromagnetic waves travelling through a vacuum, we can write

$$c = f\lambda \quad (1.2)$$

When an electromagnetic wave travels from one medium to another, its frequency remains the same, while, in general, its speed will change. This has the consequence that the wavelength also changes as an electromagnetic wave travels from one medium to another.

The identification of light with electromagnetic waves was a major development in the history of physics and was the source of much progress. However, scientists now recognize that the identification was not entirely correct. Electromagnetic waves can account for many

of the properties of light, but not all of them. For this reason, rather than saying that light is electromagnetic waves, we prefer to say that electromagnetic waves provide a model of the phenomenon of light. **The electromagnetic wave model of light** helps us to understand the behaviour of light but it is not the whole story. Other models are also useful.

The photon model of light

At present, the most complete scientific account of light involves a branch of physics called **quantum theory**. The full quantum theory of light is much too complicated to describe here, but during its development another simple model of light emerged that was quite different from the electromagnetic wave model yet, under the right circumstances, just as valuable. This alternative model is known as the **photon model of light**. According to the photon model, a ray of light of frequency f can be thought of as consisting of a stream of separate particles called **photons**. Each of these photons carries an identical amount of energy, which we can denote by the Greek letter epsilon, ε , that relates directly to the frequency of the ray, and is given by

$$\varepsilon = hf \quad (1.3)$$

The quantity h in this equation is one of the fundamental constants of physics; it is the **Planck constant** and it is given by $h = 6.626\,069 \times 10^{-34} \text{ J s}$ although $6.63 \times 10^{-34} \text{ J s}$ will do for most calculations. It is often convenient to measure the energy of photons in terms of a quantity called the **electronvolt (eV)** rather than in joules. In order to convert between joules and electronvolts, note that $1 \text{ eV} = 1.602 \times 10^{-19} \text{ J}$. For the purposes of applying Equation 1.3 to obtain the photon energy in eV rather than joules, it is also useful to note that $h = 4.14 \times 10^{-15} \text{ eV s}$.

The photon model of light is of particular importance when considering the interaction of light with atoms. Individual atoms can absorb or emit only entire photons. Thus, when a cloud of atoms is illuminated by a beam of light there is no possibility of a single atom acquiring half a photon's worth of energy directly from the beam. This is a subject to which we shall return in Section 1.3.2, when we further develop the 'background science' of light.

It is important to realize that neither the electromagnetic wave model nor the photon model should be regarded as 'true'. Light is neither a wave nor a particle but, under the appropriate conditions, it may exhibit wave-like or particle-like behaviour; both possibilities are encompassed by the quantum theory.

JAMES CLERK MAXWELL (1831–1879)

Maxwell (Figure 1.15) came from a well-to-do Scottish family. He was a shy child and his experience of school was not a happy one. Despite this, he maintained a curiosity about the natural world, and at the age of 15 he made the first of many discoveries: in this case the technique of drawing an ellipse using two pins and a thread. He went on to the University of Edinburgh and then to Trinity College, Cambridge. His interests were broad and he made significant advances in such diverse fields as colour vision, the stability of Saturn's rings, and the kinetic theory of gases. However, his outstanding achievement was the development of the mathematical theory that linked electricity and magnetism. His academic career took him to posts in Aberdeen and at King's College in London, before the death of his father in 1865 at which time he returned to the family home to become a gentleman-farmer engaging in scientific research. In 1874 he was invited to become the first Cavendish Professor of Experimental Physics at Cambridge, and was instrumental in establishing a world-class laboratory for physics at that university.



Figure 1.15 James Clerk Maxwell. (Science and Society Picture Library)

QUESTION 1.3

The overall shape of the graph in Figure 1.3b clearly indicates the phenomenon of limb darkening. Additionally, the graph includes a good deal of small-scale structure: it is not smooth but contains many tiny peaks and troughs. How can you account for this on the basis of the information given in this section?

QUESTION 1.4

The electromagnetic waves used to model the various colours of visible light have wavelengths in a vacuum in the approximate range 400 nm to 700 nm. What is the corresponding range of frequencies?

QUESTION 1.5

In terms of the photon model of light, what is the approximate range of photon energies corresponding to the range of wavelengths discussed in Question 1.4? Express your answers in joules and in electronvolts.

1.3 Seeing the Sun's inner atmosphere

1.3.1 Introducing the chromosphere

Although the majority of the Sun's light comes to us from the photosphere, we also receive small amounts of light from layers of hot, thin gaseous material that surround the photosphere. These outer layers of the Sun may be regarded as the Sun's 'atmosphere', though the term must be treated with the same degree of caution that we used when referring to the photosphere as the Sun's 'surface'.



Figure 1.16 The solar corona during a total solar eclipse. (J. Durst, Schonenberg)

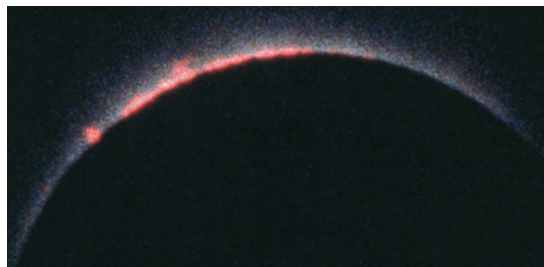


Figure 1.17 The solar chromosphere during a total solar eclipse. (G. East)

In a **white light image** of the Sun, recorded using ordinary visible light (such as Figure 1.1), the feeble light from the Sun's atmosphere is normally drowned out by the overwhelming brilliance of the photosphere. Nonetheless, under the right circumstances, it is possible to see the Sun's atmosphere, as Figure 1.16 shows. The picture was taken during a **total eclipse of the Sun** – an infrequent but predictable event that occurs somewhere on Earth on average every eighteen months or so, when the Moon passes between the Earth and the Sun, and entirely blocks the light from the photosphere for a few minutes. (It is a remarkable coincidence that, although the Sun and Moon differ greatly in size and distance from the Earth, their diameters and distances are just right to allow such a very precise blockage to occur.)

The black circle in the middle of Figure 1.16 is the silhouette of the Moon. The bright halo that surrounds it is the solar atmosphere. As you can see, the atmosphere is very extensive. For the most part it is a pearly white, but very close to the photosphere there is a narrow region with a pink or reddish tinge: this can be clearly seen in Figure 1.17. The coloured layer, which is a few thousand kilometres thick, is called the **chromosphere** (meaning the 'sphere of colour'), and constitutes the inner (or lower) solar atmosphere. The chromosphere will be our main concern in this section. The extensive outer (or

upper) solar atmosphere is called the *corona*. This will be the subject of Section 1.4.

Although eclipse studies led to the initial identification of the chromosphere, and continue to play a role in its scientific investigation, they are not the only source of chromospheric information. Fortunately, much may be learnt from observations of the full solar disc, provided they are restricted to wavelengths where the chromosphere is more prominent than the photosphere. Just such a restricted wavelength view is shown in Figure 1.18. In this particular case the image was produced by red light in a narrow range of wavelengths centred on 656.3 nm. At these wavelengths, for reasons that will be explained shortly, hydrogen atoms throughout the chromosphere are highly effective absorbers and emitters of radiation. As a result, the chromosphere absorbs most of the 656.3 nm radiation coming from the photosphere, but its own emissions at that wavelength are quite prominent. It is these emissions that are mainly responsible for the reddish hue that gives the chromosphere its name. The particular kind of restricted wavelength view shown in Figure 1.18 is called an **H α image** (pronounced ‘aitch alpha’): the H indicates that the emitted light is coming from hydrogen atoms, and the α indicates that this is the first (longest) wavelength of visible light at which hydrogen atoms have this particular effectiveness as absorbers and emitters. The other, successively shorter, visible wavelengths at which hydrogen behaves in this way are 486.1 nm, 434.0 nm and 410.1 nm, which are respectively denoted by H β , H γ and H δ (β , γ , δ are the Greek letters ‘beta’, ‘gamma’ and ‘delta’ respectively). Restricted wavelength images at other wavelengths, associated with other kinds of atom, are also of great value, particularly calcium H (396.8 nm) and calcium K (393.3 nm) images. As their names imply, these kinds of image involve light emitted from calcium atoms. (The letters H and K historically indicated the sequence of wavelengths and have nothing to do with the usual chemical symbols for hydrogen and potassium.) A calcium K image is shown in Figure 1.19.

The chromospheric H α (Figure 1.18) and calcium K (Figure 1.19) images are clearly very different from the photospheric white light image that we looked at earlier (Figure 1.1). In an H α image the chromosphere is mottled with bright specks, some of which are gathered together into extensive bright regions called **plages** (the French word for ‘beach’ that is pronounced ‘plah-je’). These are often seen in parts of the chromosphere that are directly above the active regions of the photosphere that contain sunspots and are especially prominent in calcium K images. Visible in the bottom half of the H α image (Figure 1.18) is a long, winding dark feature called a **filament**. Filaments are quite common in H α images; they are caused by huge clouds of relatively cool gas held high above the

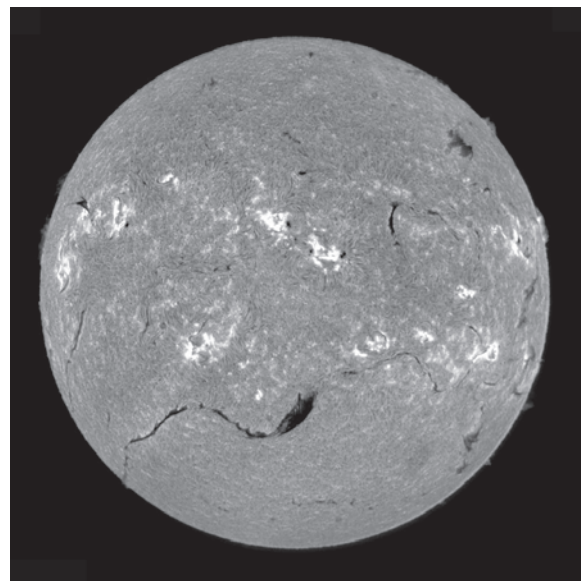


Figure 1.18 An H α image of the Sun, produced by light from a narrow range of wavelengths centred on 656.3 nm. (Big Bear Solar Observatory)

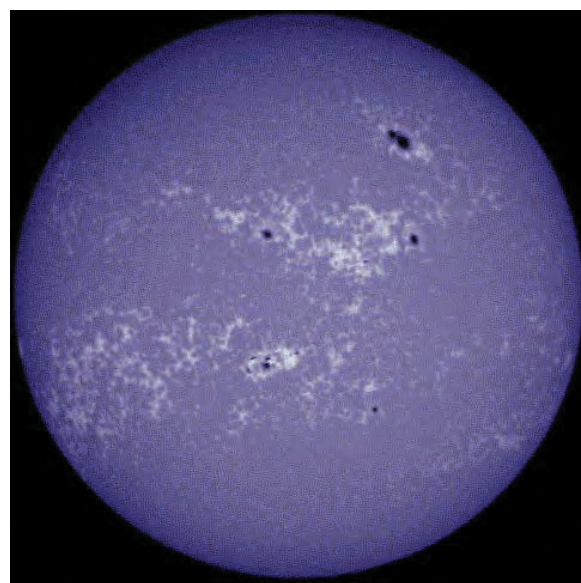


Figure 1.19 A calcium K image of the Sun. The image shows emission in a narrow range of wavelengths centred on 393.3 nm. As in the H α image (Figure 1.18), this image reveals the structure of gas in the chromosphere. (Marshall Space Flight Center/NASA)

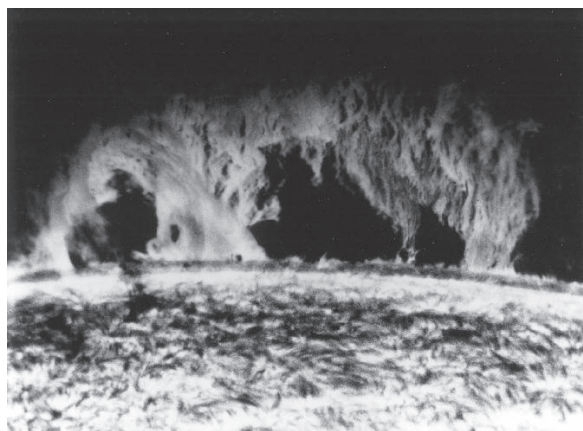


Figure 1.20 An $H\alpha$ solar prominence on the limb of the Sun. (Big Bear Solar Observatory)

chromosphere by magnetic forces. These are the same clouds that account for the **prominences**; such as the one seen above the limb of the Sun in Figure 1.20. Overall, the chromosphere is much less uniform than the photosphere. One solar physicist has even described the chromosphere as ‘a layer of froth stirred up by the photosphere’.

To get a better understanding of the chromosphere, we really need to gain some insight into the processes that give rise to $H\alpha$ images and to the other, similar, images obtained at different wavelengths. Only with the aid of such insight can the radiation emitted by the chromosphere be properly interpreted and used as a source of information about the physical conditions of the Sun, such as temperature, pressure and structure. The starting point for this kind of investigation is the technique of *spectroscopy*.

1.3.2 Spectroscopy and sources of light

Most sources of light emit over a range of wavelengths. (The one well-known exception to this is the laser, which is a device that produces light at a single wavelength.) Such a range of wavelengths is commonly called a **spectrum** (plural *spectra*). **Spectroscopy** concerns the production and study of spectra.

Continuous spectra

Many sources of light emit over an unbroken range of wavelengths. Such sources are therefore said to have **continuous spectra**.

If a narrow beam of light passes through a glass prism, the beam will split up in such a way that different wavelengths travel in different directions. (This process is shown in Figure 1.21.) If the original beam contained just a few well separated wavelengths the result would be a set of quite separate and distinct images, each with its own characteristic colour (wavelength). However, if the beam came from a source which produces a continuous spectrum it would typically contain all visible wavelengths, and the result of passing it through the prism would be a multicoloured band somewhat similar to a rainbow.

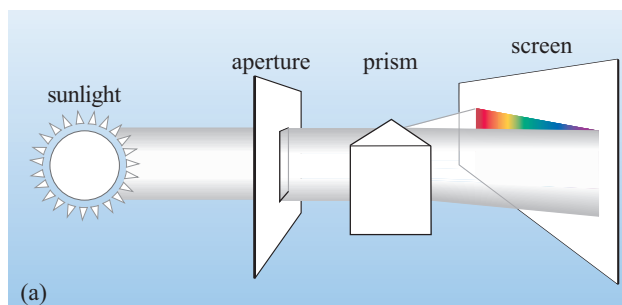
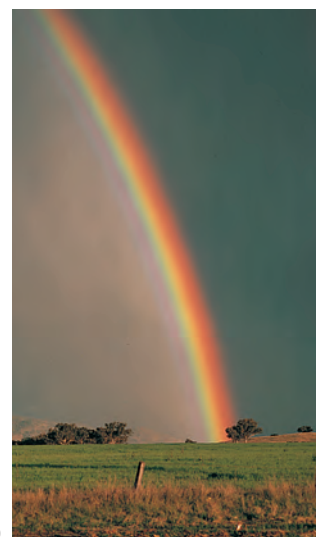


Figure 1.21 (a) The formation of a continuous spectrum by use of a glass prism. (b) A naturally occurring spectrum – a rainbow. Here the light from the Sun is dispersed by passing through raindrops. ((b) G. Garrad/Science Photo Library)



In the past, scientists engaged in spectroscopy spent a good deal of their time examining the kinds of multicoloured band described above. Nowadays things are different in that it is more common for spectral information to be given in the form of a graph. When spectra are presented in this way the horizontal axis of the graph usually shows wavelength or frequency (or sometimes photon energy). The vertical axis of the graph normally indicates the ‘brightness’ of the spectrum at any given wavelength. Figure 1.22 shows the relationship between this graphical representation of a spectrum and the coloured band spectrum. Unfortunately, the quantities used to measure spectral brightness have rather complicated definitions. In this chapter the vertical axes of some graphical spectra are labelled **spectral flux density**. At any given wavelength λ , the spectral flux density, F_λ , can be determined by the following procedure.

- Using an appropriate detector of area 1 m^2 , pointed directly towards the source, measure the rate at which energy from the source is delivered to the detector by electromagnetic waves with wavelengths in a fixed narrow range, $\Delta\lambda$, centred on λ .
- Divide the measured rate of energy detection by the wavelength range $\Delta\lambda$ to obtain the detected power per square metre per unit wavelength range, typically measured in units of $\text{W m}^{-2} \mu\text{m}^{-1}$ or in $\text{W m}^{-2} \text{nm}^{-1}$. This is the value of F_λ at wavelength λ .

In fact, most of the graphical spectra in this book will have vertical axes that show **relative spectral flux density**. In such cases, the spectral flux density at any wavelength is expressed as a *fraction* of some arbitrarily chosen reference value and there will be no SI units shown on the axis.

The black-body spectrum

There are several different types of physical process that can give rise to continuous spectra. One particularly important class of continuous spectrum is called the **black-body spectrum** (or sometimes a Planck spectrum, after the German physicist Max Planck (1858–1947)).

There are two key features of sources that produce black-body spectra. The first is that the emission of light arises as a result of the material that makes up the source being at a relatively high temperature. From everyday experience, you can probably think of examples that support the statement that ‘hot things glow’. In such a source, the energy that is emitted as light, has its origin in the internal, or thermal, energy of the material that makes up the source. Not surprisingly, such sources are termed **thermal sources** of radiation. It is important to appreciate that not all sources of light are thermal in origin; those sources in which the energy emitted as light does not derive from the thermal energy of the source would be called **non-thermal sources** of radiation. An everyday example of a non-thermal source of light is a television screen, which produces a copious amount of light but is not hot.

The fact that a source of light is thermal is not a sufficient condition for it to generate a black-body spectrum; indeed, thermal sources need not even produce continuous spectra. To generate a black-body spectrum, a second condition must be met in that light within the source is much more likely to interact with the material of the source than to escape. This means that light generated within the

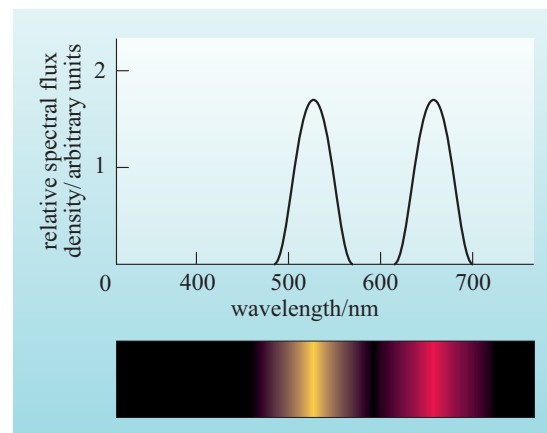


Figure 1.22 The relationship between the band spectrum and a graphical representation of a spectrum.

source is in some sense ‘trapped’, and is only likely to escape after there has been considerable interaction with the material within the source. Therefore, a common feature of sources that approximate well to being black-body sources is that they are opaque. This is because it is unlikely that a photon could cross the material that makes up the source without undergoing an interaction.

- Why is a source of light that is transparent unlikely to form a black-body spectrum?
- If the material that is emitting light is transparent then it is more likely that photons that are generated would escape from the source rather than interact with the material of the source. Hence the source is unlikely to produce a black-body spectrum.

An everyday example that satisfies the conditions for forming a black-body spectrum reasonably well is the tungsten filament of a light bulb. The filament is heated and light is produced not just at the surface, but throughout the bulk of the filament. Because solid metals are far from being transparent, most of the light that is produced by the filament does not escape to be seen, but interacts with the material of the source. Near the surface of the filament however, the assumption that light is much more likely to interact with the material of the source than to escape starts to break down, and this results in the spectrum only being an approximation to the ideal black-body conditions.

The concept of the black-body spectrum is useful because many astronomical sources produce continuous spectra that are a reasonably good approximation to the black-body form. However, it should also be noted that there are several other types of physical mechanism that can produce continuous spectra.

Figure 1.23 shows a set of black-body spectra from sources at temperatures between 3000 K and 6000 K. Note that the shape of the curve varies with temperature, but that the form of the curve is relatively simple: it has just one broad maximum or peak.

QUESTION 1.6

Figure 1.23 clearly indicates that the observed spectral flux density from the hottest source is greater than that from the coolest. Is it necessarily the case that in a fixed wavelength range more energy will arrive per second from any black-body source at 6000 K than from *any* black-body source at 3000 K?

In any graph that shows how relative spectral flux density varies with wavelength for a black-body source, the height of the graph will depend on factors such as the size and distance of the source, and the selected reference value. However, in all such graphs the overall shape of the curve is solely determined by the *temperature* of the source. This means, in particular, that the peak of each curve occurs at a wavelength, λ_{peak} , that characterizes the source’s temperature, irrespective of the height of the peak. In fact, there is a simple law, called **Wien’s displacement law**, that relates the value of λ_{peak} to the temperature, T , of the source:

$$(\lambda_{\text{peak}}/\text{m}) = \frac{2.90 \times 10^{-3}}{(T/\text{K})} \quad (1.4)$$

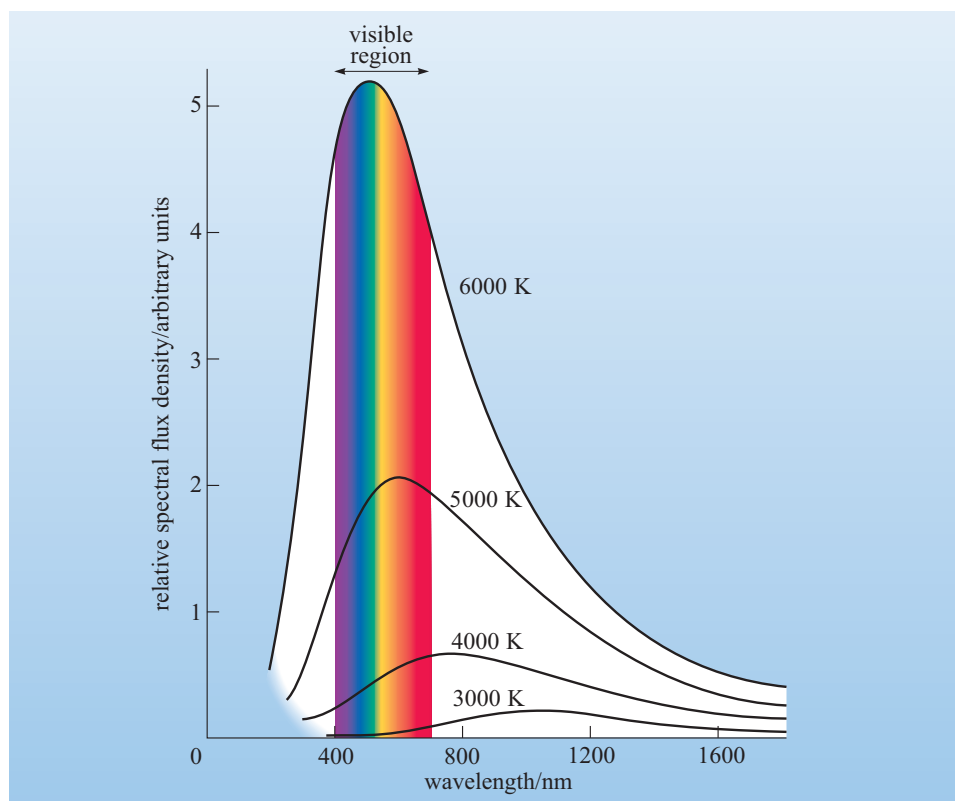


Figure 1.23 Graphical spectra for black-body sources at temperatures between 3000 K and 6000 K. The black-body sources all have the same surface area and are at the same distance from the detector. (Note that the spectra extend well beyond those wavelengths that correspond to the various colours of visible light. More will be said about these other emissions in Section 1.4.)

With the aid of Wien's displacement law it is a simple matter to determine the temperature of any source of light, *provided that it is a black-body source*. Such sources are not common, but many real sources, including the Sun and other stars, are reasonable approximations to black-body sources, so their temperatures may be estimated by this spectral technique.

QUESTION 1.7

If you were to heat a metal ball, so that its temperature steadily increased, you would find that above a certain temperature the ball would start to emit a dull red glow. As the temperature increased further the ball would become brighter and the colour would change from red to orange-white to yellowish-white to white. How would you explain these changes in appearance?

The shapes of the curves that describe the spectra of black-body sources are of great importance in science. Such curves are usually referred to as thermal radiation curves; or **Planck curves** or **black-body radiation curves**. While the black-body spectrum is important, it should also be noted that other physical processes may also produce continuous spectra, but those spectra will generally have a different shape from those of black-body sources.

Line spectra: absorption and emission

If a beam of light from a black-body source passes through a thin (low-density) gas of atoms, the spectrum of the emerging beam will generally include a number of narrow dark lines. These lines are called **absorption lines** and correspond to narrow ranges of wavelength that have been wholly or partly absorbed by atoms in

the gas (a process described in Box 1.2). This situation is illustrated in Figure 1.24, which includes a graph of the so-called **absorption spectrum** that arises.

If, instead of the emerging beam, the light emitted by the gas itself is examined, it will be found that its spectrum consists of a number of narrow *bright* lines. These lines are called **emission lines**, and a spectrum composed of them is called an **emission spectrum**. For many gases, the bright lines emitted cover the same narrow wavelength ranges as the dark lines in the absorption spectrum. The graph of such an emission spectrum is also included in Figure 1.24.

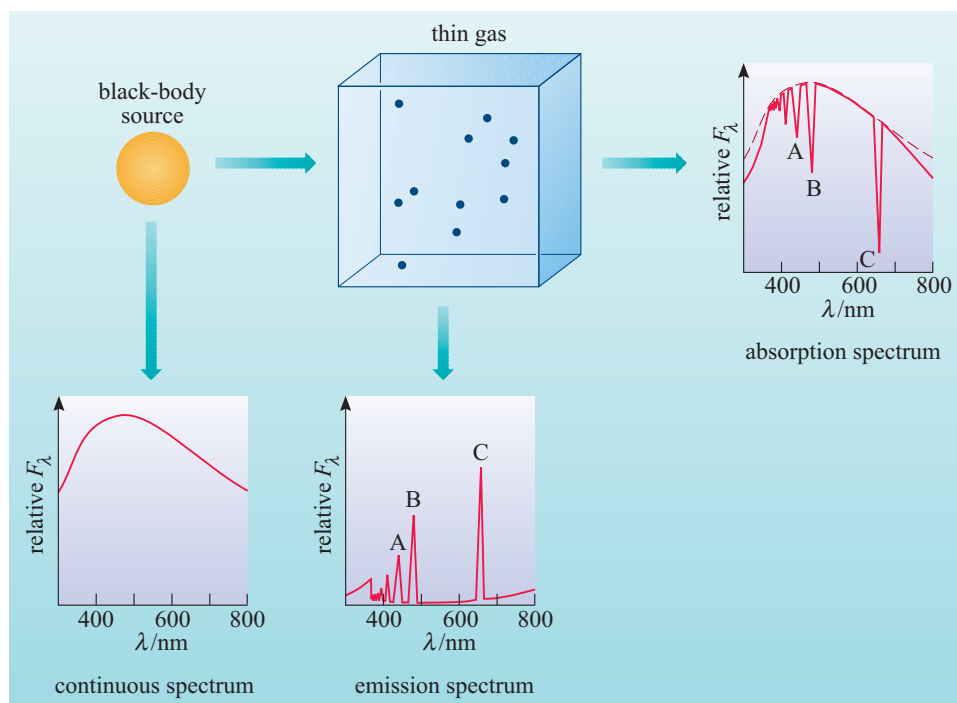


Figure 1.24 Three kinds of spectrum – continuous, absorption and emission – seen by observing a black-body source and a thin gas of atoms from various directions. The dashed line in the absorption spectrum shows the continuous spectrum that would have been observed in the absence of the gas.

Emission spectra and absorption spectra are both examples of **line spectra**.

Figure 1.25 shows the line spectra produced by a number of different gases. The occurrence of such spectra is a consequence of the fact that each of the gases is composed of a characteristic type of atom that has its own internal structure. (It is also relevant that the gas is sufficiently thin to ensure that the atoms do not significantly influence one another.)

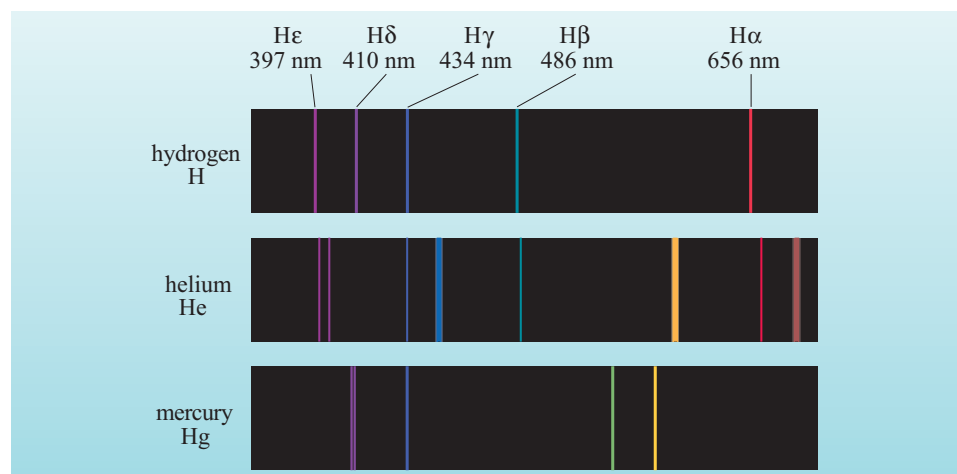


Figure 1.25 The line spectra produced by hydrogen, helium, and mercury. Note that in all cases the spectra are produced by the elements in gaseous form.

BOX 1.2 SPECTRAL LINES AND ENERGY-LEVEL DIAGRAMS

A typical atom is about 2×10^{-10} m in diameter and consists of a tiny, dense, positively charged *nucleus* surrounded by one or more negatively charged electrons. According to quantum theory, each of the electrons belonging to a particular atom may be in any of a number of allowed states, each of which is associated with some fixed amount of energy. When an electron occupies a particular state in a particular atom, the atom has the energy associated with that state. Thus, changes in the pattern of occupied states within an atom entail changes in the total amount of energy possessed by the atom. A diagram showing the energy associated with each of the allowed states in a particular kind of atom is called an **energy-level diagram**. The simplest energy-level diagram, that of a hydrogen atom (which has just one electron), is shown in Figure 1.26.

When a hydrogen atom has the lowest energy level, labelled E_1 , it is said to be in its **ground state**. Above this energy, when the atom has energy E_2 , E_3 , etc., it is said to be in an **excited state**.

To see how the behaviour of atoms can account for the existence of line spectra, consider an atom which initially occupies a state of energy E_i (the subscript i reminds us that this is the initial energy). Such an electron may make a transition to some other state of higher energy E_f (f for the final, higher energy state), provided that the atom can acquire the necessary additional energy, $E_f - E_i$. If the atom is bathed in light from a source that produces a continuous spectrum, as in Figure 1.24, the requisite energy can be absorbed directly from the light. Now, you might think that energy absorption of this kind would involve a wide and continuous range of wavelengths, but that is not the case. The absorption of light by an atom is one of those processes that is best described by the photon model of light discussed in Box 1.1. The energy is acquired in one gulp, as it were, by the absorption of a single photon of just the right energy $\varepsilon_{fi} = E_f - E_i$. It follows from Equation 1.3 that such a photon

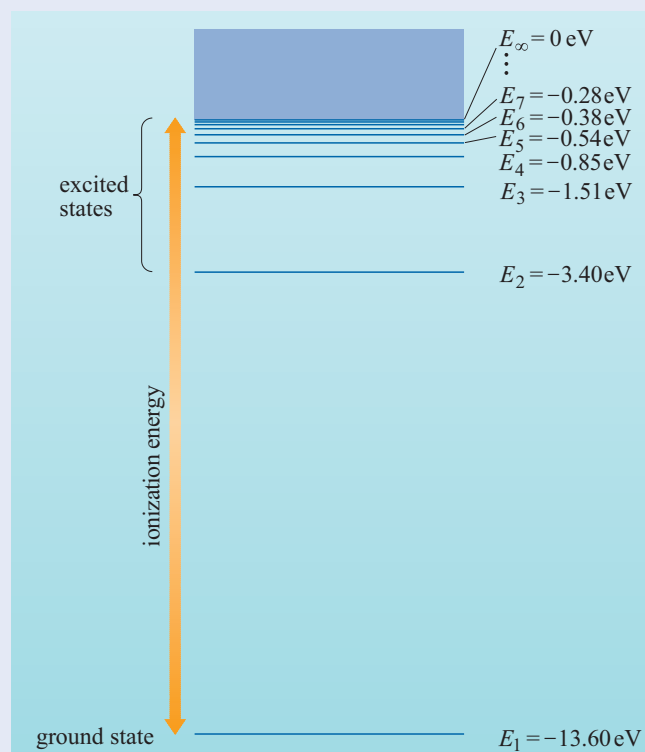


Figure 1.26 The energy-level diagram of a hydrogen atom. There are infinitely many energy levels but those of highest energy are too closely crowded together to be shown separately on a diagram of this kind. Note that the energies are given in terms of electronvolts ($1 \text{ eV} = 1.602 \times 10^{-19} \text{ J}$).

corresponds to electromagnetic radiation of a specific frequency, namely

$$f_{fi} = \frac{1}{h}(E_f - E_i) \quad \text{absorption} \quad (1.5)$$

There is a limit to the amount of energy that a hydrogen atom can absorb without the electron and proton being split apart. This limit is known as the **ionization energy** and for the hydrogen atom it has a value of 13.60 eV.

QUESTION 1.8

Write down a formula for the corresponding wavelength, λ_{fi} , of the radiation that must be absorbed if a transition requiring an energy increase $E_f - E_i$ is to take place.

Applying the result of Question 1.8 to the gas of atoms shown in Figure 1.24, it should be clear that if a sufficient number of atoms in the gas increase their energy by $E_f - E_i$ every second, then a good deal of incoming radiation at the corresponding wavelength λ_{fi} will be absorbed, with the result that the spectrum of the emerging beam will have a dark absorption line at that wavelength. Since there are many possible values of E_i and E_f for any particular kind of atom, it follows that there are many possible values of $(E_f - E_i)$ and hence many, quite distinct wavelengths at which absorption lines might occur. (Whether a particular line is seen or not depends on the *rate* at which the relevant transition is occurring in the gas.) Furthermore, since each kind of atom has its own characteristic energy-level diagram, different gases will produce different absorption spectra – just as you saw in Figure 1.25.

The above account of absorption spectra requires a large number of atoms to absorb energy every second, at many different wavelengths. What happens to all this absorbed energy? The answer is that most of it is very quickly emitted again, and it is this emission that accounts for the emission spectrum that an illuminated gas produces. Once an electron occupies a state associated with an energy E_f that is higher than the energy associated with one or more other states, it may be possible for the electron to make a spontaneous transition to one of those states of lower energy. In particular, if the electron originally came from a state of energy E_i , one of many possibilities is for it to return to that same state. Under these circumstances the atom would have to shed an amount of energy $\varepsilon_{fi} = E_f - E_i$. Once again, quantum theory demands that if the energy released in a transition of this kind is given off as light then it must take the form of a single photon corresponding to electromagnetic radiation of frequency

$$f_{fi} = \frac{1}{h}(E_f - E_i) \quad \text{emission}$$

A gas is therefore capable of producing emission lines at exactly the same wavelengths that it produces absorption lines. However, and this is a crucial point, the emission is equally likely to occur in any direction. Thus, energy absorbed from the incoming beam in Figure 1.24 will not be entirely replaced by emitted energy; rather emitted energy will be given off in all directions, allowing the emission spectrum to be seen from many different angles, whereas the absorption spectrum can be seen only by looking into the beam.

1.3.3 Solar spectroscopy and the structure of the chromosphere

A good deal of what is known about the Sun has been learnt by studying its spectrum. A photograph of the visible spectrum is given in Figure 1.27. As you can see it is essentially an absorption spectrum; the coloured bands are crossed by a large number of narrow absorption lines corresponding to various atomic transitions. In fact, something like 25 000 lines have been identified in the visible region; many originate in the photosphere, which is also responsible for nearly all the light between the lines, but some of the lines carry information about the chromosphere. Unfortunately, owing to the problems of reproducing colour images, not all of the features of Figure 1.27 are as clear as they could be, so a sharper

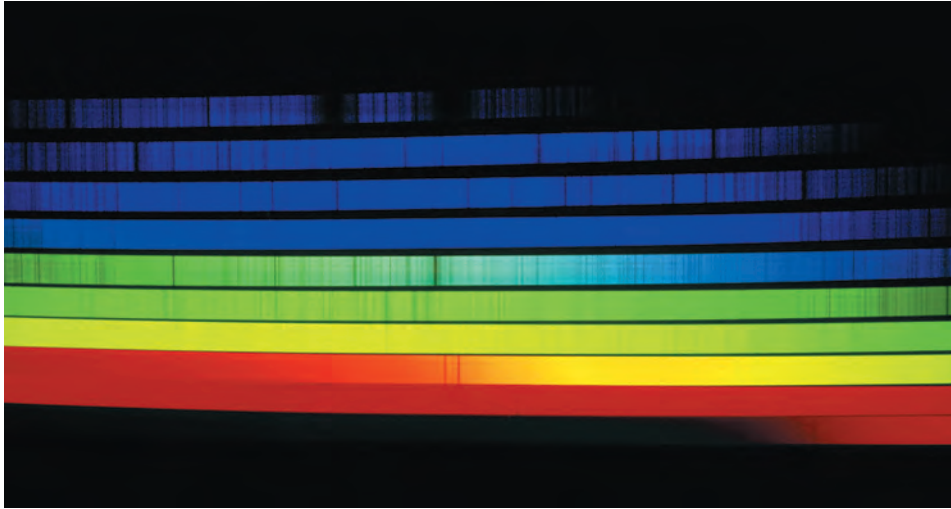
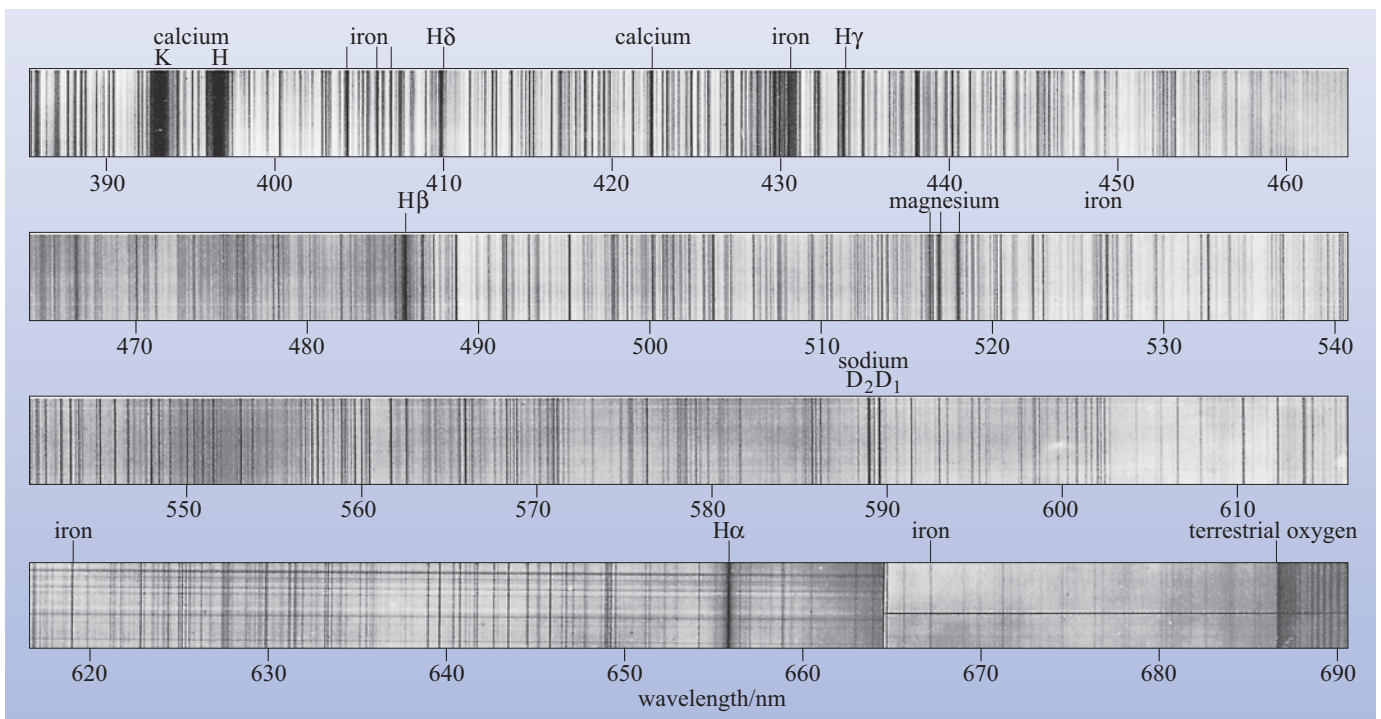


Figure 1.27 The solar absorption spectrum. Note that for convenience of display, the spectrum has been cut into sections and consecutive sections have been stacked vertically in sequence. (National Solar Observatory)



black and white image of the Sun's spectrum is given in Figure 1.28. Some of the darkest lines in this image have been labelled and a wavelength scale (in nanometres) has been included to help you locate them. It is worth noting that even the darkest lines are not completely black; they are simply darker than their surroundings. Indeed, it is often the faint emissions found within the darkest lines that are most informative about the chromosphere. A particularly strong absorption line of this sort can be seen around 656.3 nm, that is the wavelength of the $H\alpha$ image we looked at earlier. Not surprisingly, this is called the **$H\alpha$ absorption line**. Another prominent absorption line due to hydrogen ($H\beta$) at 486.1 nm, and two lines due to calcium at 396.8 nm and 393.3 nm, were all mentioned earlier as important wavelengths for studies of the chromosphere.

Figure 1.28 A black and white image of the solar spectrum. Note that for convenience of display, the spectrum has been cut into sections and consecutive sections have been stacked vertically in sequence. (The horizontal streaks on the spectra are artefacts.) (Kitt Peak National Observatory)



Figure 1.29 Sir (Joseph) Norman Lockyer (1836–1920). Lockyer developed an interest in astronomy when he was working as a civil servant, and he went on to become a pioneer of solar astrophysics. His contribution to science was much wider than his chosen specialism: he founded the journal *Nature*, which he edited for fifty years. (Royal Astronomical Society)

QUESTION 1.9

Why is it not surprising that studies of chromospheric emission lines often involve wavelengths that correspond to dark absorption lines in the solar spectrum?

Chromospheric emission lines can be observed in particular detail during total solar eclipses, when the light of the photosphere is blocked by the Moon. One of these emission lines, at 587.6 nm, has a particularly interesting history. When it was first observed, during the eclipse of 18 August 1868, its origin could not be explained in terms of any of the chemical elements then known. This led Sir Norman Lockyer (1836–1920) (Figure 1.29) to propose the existence of a new element, which he named helium after the Greek word for the Sun, *helios*. Lockyer's interpretation was correct, but confirmation did not come until 1895, when the British scientist Sir William Ramsay (1852–1916) showed that helium exists on Earth.

By comparing absorption and emission lines of solar origin with lines observed in terrestrial laboratories, it is possible to determine which kinds of atom are present in the visible parts of the Sun. In this way the presence of more than 65 of the 100 or so known elements has been established. This is a notable achievement, but even more can be accomplished by examining the details of the lines – a task best done with the aid of the graphical form of the spectrum.

Figure 1.30 is a graphical representation of the solar spectrum. The dips corresponding to some of the more prominent absorption lines can be discerned. In order to explain the detailed appearance of this kind of graph it is not enough simply to identify which atoms are represented; the relative depths of the various lines must also be explained. Line depths depend on a number of factors, including the relative abundance of each kind of atom, the inherent likelihood of each atomic transition, and the proportion of atoms of a given kind that have an electron in the appropriate initial state to give rise to a particular line. This last factor will depend sensitively on temperature because, generally speaking, the higher the temperature the more likely it is that states corresponding to higher energies will be occupied. Thus, solar absorption lines provide information about chemical abundance, temperature and related quantities such as density and pressure. Similar comments apply to the chromospheric emission lines; in fact, these lines provide some of the most direct information about the composition and structure of the chromosphere.

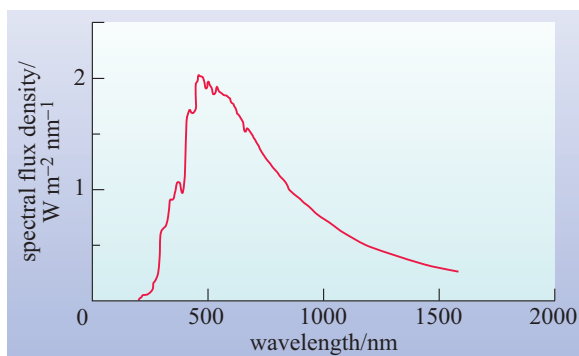


Figure 1.30 A graphical representation of the Sun's spectrum, as measured from above the Earth's atmosphere. (Phillips, 1992, from the original data of Labs and Neckel, 1968)

Information gained from spectral studies shows that the chromosphere is composed mainly of hydrogen and helium. The evidence indicates that there are between 40 and 100 helium atoms for every 1000 hydrogen atoms. (Other kinds of

study, to be discussed in Section 2.2, imply that the correct value is probably about 87, corresponding to about 25% of the mass.) In view of the prominence of the calcium lines, you might also expect that calcium atoms would be abundant, but this turns out not to be the case. The relative strength of the calcium lines results from the combination of two factors: chromospheric conditions are just right to ensure that a comparatively high proportion of calcium atoms have electrons in the correct initial states to produce those lines, and the relevant transitions have a relatively high likelihood of occurring, given electrons in the correct initial states. In fact, there are probably no more than three calcium atoms for every *million* hydrogen atoms.

As far as temperature and density are concerned, it has already been emphasized that these quantities vary with height and from one region of the chromosphere to another. Nonetheless, typical values for a ‘quiet’ part of the chromosphere are shown in Figure 1.31. The rapid increase of temperature with height is very clear. Note that the height is measured from the usual reference level at the base of the photosphere (roughly, the greatest visible depth), so only data pertaining to heights greater than about 500 km are truly ‘chromospheric’. It is thought that the chromosphere is heated mainly from below (by energy coming from the photosphere) so it is interesting, and not a little perplexing, that the temperature should rise with increasing height. Even more astonishing is the enormous rate at which the temperature rises, especially towards the top of the chromosphere. This surely indicates a very high temperature indeed for the Sun’s outer atmosphere – the corona. The high temperature of the corona and upper chromosphere, its cause and its influence will be our major concern in the next section.

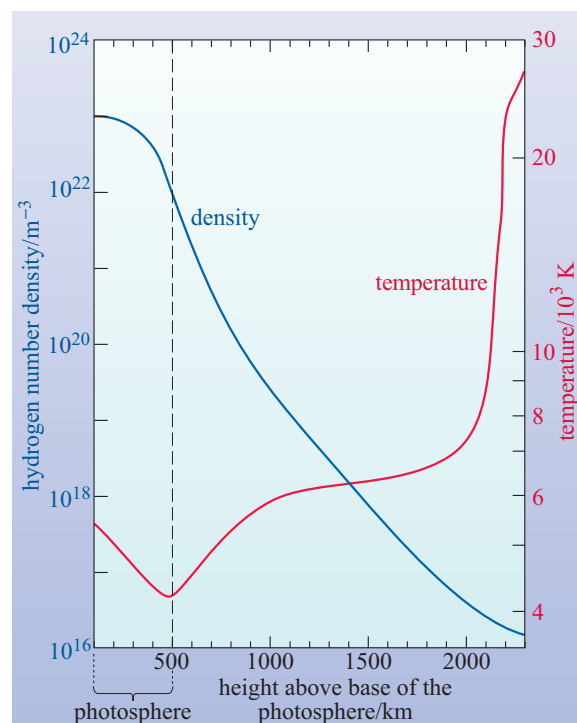


Figure 1.31 Variation of temperature and hydrogen number density (i.e. number of hydrogen atoms per cubic metre) with height throughout a ‘quiet’ part of the chromosphere. For comparison, at sea-level the Earth’s atmosphere has on average a temperature of about 288 K and a number density (of oxygen and nitrogen molecules) of about $2 \times 10^{25} \text{ m}^{-3}$. Note that this figure uses *logarithmic scales* on the vertical axes.

QUESTION 1.10

Using a certain detector it has been found that a particular source of electromagnetic waves produces a black-body spectrum with a temperature of 6000 K. Use Figure 1.23 to find the ratio of the detector reading at 400 nm to that at 700 nm. What would the ratio have been if the spectrum had been that of a black body at 5000 K?

QUESTION 1.11

Treating the Sun as a good approximation to a black body, estimate its temperature from the graphical spectrum of Figure 1.30.

1.4 Seeing the Sun's outer atmosphere

1.4.1 Introducing the corona

As you saw at the end of the last section, the temperature rises steeply in the upper parts of the chromosphere and the number of hydrogen atoms per cubic metre declines. These trends continue through a narrow and highly irregular **transition region** and on into the **corona** – the Sun's outer atmosphere.

The corona is very extensive, very tenuous, and has an extremely high temperature.

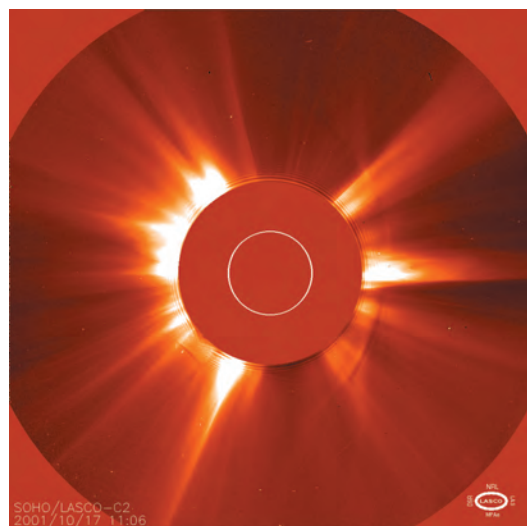
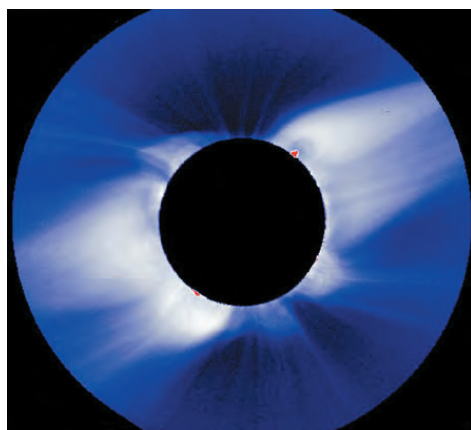


Figure 1.32 The solar corona as observed using a coronagraph on a space-based solar telescope. The white circle represents the size and location of the solar disc. (SOHO (ESA and NASA))

The extent of the corona can be gauged from photographs taken during eclipses or from measurements made with a **coronagraph** – a special kind of telescope that uses an opaque disc to block the light from the photosphere and thus produce a sort of artificial eclipse. (Building a coronagraph that works is much harder than it sounds.) High quality images, such as that in Figure 1.32 (which was taken from a coronagraph on a space-based solar observatory), show the corona stretching out to several times the radius of the photosphere. They also reveal detailed structures, often in the form of arches or rays. These features change with time, sometimes quite rapidly, and respond to the general level of (sunspot) activity visible on the photosphere. At times of low activity the corona is usually rather quiescent and elongated at the Sun's equator (Figure 1.33a); at times of high activity it is much more lively, with streamers jutting out in all directions (Figure 1.33b).



(a)



(b)

Figure 1.33 The solar corona during an eclipse at a time of (a) low, and (b) high solar activity. (NCAR)

In comparison to the photospheric emission, the total amount of light produced by the corona is small. Even the brightest parts of the corona are almost a million times less luminous at visible wavelengths than an equal area of the photosphere. Most of the observed coronal light is simply white photospheric light scattered by particles (mainly electrons) in the corona – that's why the light has a characteristic pearly white colour. In the outer parts of the corona, beyond about two solar radii from the centre of the Sun, it is even possible to see the photospheric and chromospheric absorption lines in the coronal spectrum. However, in addition to features attributable to the photosphere and chromosphere, the coronal spectrum also includes some emission lines that originate in the corona itself. The strongest of these is a green line at 530.3 nm, but there is also a prominent yellow line at

569.4 nm and a red line at 637.4 nm. An image of the corona in the light of the green 530.3 nm line is shown in Figure 1.34.

When the strongest of these coronal emission lines was discovered during the eclipse of 7 August 1869, its origin was unknown. It was suggested that it might be due to a previously undiscovered element, which was tentatively named coronium. But, in contrast to the story of helium, no other evidence was found to support this hypothesis, and the cause of this and other coronal lines remained one of the most challenging problems in solar spectroscopy for over 70 years. It was finally solved in the early 1940s by the Swedish astrophysicist Bengt Edlén (1906–1993). In a series of experiments Edlén proved that an ionized iron atom that has lost half of its normal complement of 26 electrons has its energy levels altered in such a way that there is a transition capable of producing 530.3 nm radiation (no such transition exists in a neutral iron atom that retains all 26 electrons). Since neutral iron (chemical symbol, Fe) has 26 electrons in each of its atoms, it follows that iron atoms that have lost half their electrons will have a net positive charge that is equal in magnitude but opposite in sign to the charge of the 13 lost electrons. Such atoms, or, to be more precise, such ions are conventionally denoted by the symbol Fe^{13+} . Thus, Edlén had established that the green coronal emission was due to Fe^{13+} . The other coronal emission lines known at the time were also shown to be due to highly charged ions.

The presence of ions in the Sun is not surprising. Temperatures throughout the Sun are so high that most of the atoms are ionized to some extent. Some of the effects due to the presence of these ions are already familiar to you: the prominent H and K lines due to calcium (chemical symbol, Ca) in the chromosphere come from Ca^+ ions. In addition, much of the light from the photosphere is emitted when free electrons, liberated when atoms are ionized, combine temporarily with neutral hydrogen atoms to produce H^- ions. However, the discovery that a significant amount of Fe^{13+} exists in the corona, was a surprise. The presence of such highly ionized atoms implied that the temperature of the corona was very high indeed – at least 10^6 K. Previously it had been thought that temperatures decreased as distance from the photosphere increased; the discovery that this was not the case replaced the enigma of the emission lines by the mystery of the mechanism that could be responsible for such high temperatures.

In part, the very high temperature of the corona is due to a rather peculiar property of gases at low densities and high temperatures. Over a range of temperatures between about 10^5 K and 10^6 K, the rate at which coronal gas would lose energy by radiating electromagnetic waves is lower than at the temperatures at either end of this range (i.e. at about 10^5 K or 10^6 K). This is quite different to our everyday experience of how hot material behaves, in that we usually expect, and usually find, that the rate at which a body loses energy simply increases with temperature. Coronal gas that is at a temperature between 10^5 K and 10^6 K is in an unstable state: if the gas is heated such that its temperature exceeds 10^5 K, then the gas is unable to cool efficiently and its temperature quickly rises to a temperature in excess of 10^6 K. However, this instability only explains why the temperature is so high; it does not explain the source of the heating for the corona.

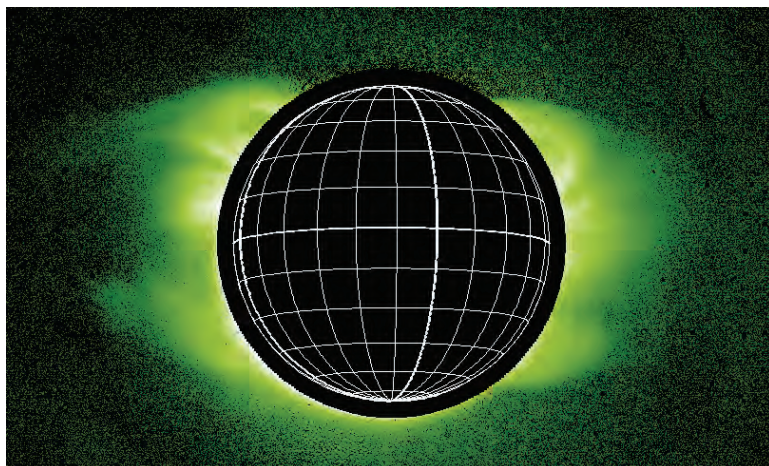


Figure 1.34 An image of the corona taken at a wavelength of 530.3 nm. (SOHO (ESA and NASA))

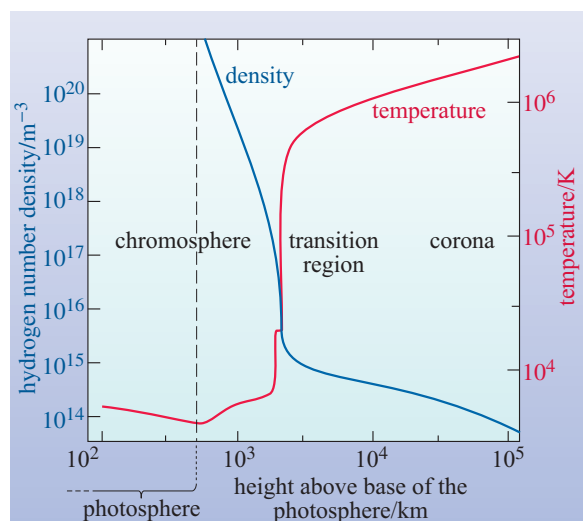


Figure 1.35 Variation of temperature (red) and hydrogen number density (blue) with height above the base of the photosphere. Note that all the scales are logarithmic. This figure incorporates the information shown in Figure 1.31, but extends to show how temperature and density vary from the photosphere to the high regions of the corona. (Adapted from Gabriel, 1976)

The corona cannot be heated simply by energy radiated by the photosphere or by heat conducted through the chromosphere; a basic law of physics (the **second law of thermodynamics**) prevents the transfer of energy from a cooler body to a hotter one by either of these methods. Nonetheless, there can be little doubt that the necessary energy does come from the lower regions of the Sun; the question is, how? It seems that the behaviour of the magnetic field in the lower parts of the corona plays an important role in heating the coronal gas. However the precise mechanism is still elusive and the absence of a universally agreed solution to the problem of coronal heating is not due to a lack of ideas but is indicative of the extreme difficulty of producing detailed theoretical descriptions of such complex regions.

Like the chromosphere, the corona is highly non-uniform and conditions vary enormously from place to place within it. Nonetheless it is possible, using spectroscopic information and basic physical principles, to obtain a good idea of ‘average’ conditions in a quiet part of the corona. Figure 1.35 shows the variation of temperature and hydrogen number density with height under such quiet conditions in the chromosphere and the inner part of the corona. The graphs are essentially extensions into the corona of the chromospheric data shown in Figure 1.31. Note that the location of the transition region separating the chromosphere

and the corona is defined by the very rapid temperature rise just below 2500 km. The precise location depends on local conditions; it might well be as high as 8000 km on some occasions. Also note that, at the greatest heights shown in Figure 1.35, the temperature of the corona is still increasing. Temperatures of 3×10^6 to 4×10^6 K are not uncommon, and even higher temperatures are sometimes attained over limited regions.

You might think that such enormous temperatures would make the corona a powerful source of electromagnetic radiation, but this is not so. The density of the corona is very low, and such a relatively small amount of matter, even at a very high temperature, is a very poor emitter compared with the cooler but much denser photosphere. Nevertheless, the high temperature of the corona does have important implications for the electromagnetic radiation that is emitted.

- Why is the spectrum of the corona unlikely to be a black-body source of radiation?
- The corona is a very low-density region that is far from being opaque. Sources that produce spectra that are a good approximation to black-body spectra tend to be highly opaque. Hence it is unlikely that the corona would exhibit a black-body spectrum.

Despite the fact that the corona is *not* a black-body source of radiation, it is still possible to estimate the wavelength at which it will radiate. A rule of thumb that can be applied to *any* thermal source is that if the source is at a temperature T , then the typical photon energy ε that would be emitted by the source is given by

$$\varepsilon \sim kT \quad (1.6)$$

where k is the **Boltzmann constant**, which has a value of $1.38 \times 10^{-23} \text{ J K}^{-1}$. The symbol ‘ \sim ’, which is often called ‘twiddles’, is used to express rather approximate relationships – the answer may be correct only to within a factor of ten or so. In applying Equation 1.6 to a real problem, the equation is evaluated as if the ‘ \sim ’ were an equals sign, but it must be remembered that the answer is only an approximate solution. There are no hard and fast rules about how to quote the numerical answers from this sort of equation: while it is usual practice to give results to one significant figure, it must be borne in mind that the implied level of precision is often not justified. Equation 1.6 is very approximate because it refers to *any* thermal source, not just the special case of black-body emission. In general, a thermal source will produce a spectrum that is more complex than the black-body form; it may, for instance, appear to be a mixture of a continuous spectrum and a line spectrum. Equation 1.6 allows us to estimate, within a factor of ten or so, the typical photon energies that are likely to be observed in such a spectrum.

Let’s return to the solar corona with an example of the use of Equation 1.6 that will allow us to estimate the wavelength of its thermal emission.

EXAMPLE 1.1

Assuming that the solar corona is a thermal source with a temperature of $2 \times 10^6 \text{ K}$, calculate (a) the typical energy (in eV), and (b) the corresponding wavelength of photons that would be emitted.

SOLUTION

(a) The typical photon energy can be found by applying Equation 1.6, using a value of $T = 2 \times 10^6 \text{ K}$,

$$\varepsilon \sim kT$$

$$\varepsilon \sim (1.38 \times 10^{-23} \text{ J K}^{-1}) \times (2 \times 10^6 \text{ K}) = 2.76 \times 10^{-17} \text{ J}$$

which, in terms of electronvolts, is

$$\varepsilon \sim (2.76 \times 10^{-17} \text{ J}) / (1.602 \times 10^{-19} \text{ J eV}^{-1}) = 172 \text{ eV}$$

Since the symbol \sim implies a very approximate relationship, the answer should be quoted to just one significant figure. Hence the typical photon energy is $2 \times 10^2 \text{ eV}$.

(b) The wavelength that is typical of emission from the corona can be found from the photon energy ε . The first step is to write an equation for λ in terms of ε . Equation 1.2 can be rearranged to make λ the subject,

$$\lambda = c/f$$

and from Equation 1.3,

$$f = \varepsilon/h$$

Combining these two equations gives the required equation,

$$\lambda = hc/\varepsilon$$

Now the wavelength can be calculated using the value of ε obtained in part (a). (It does not matter whether the value used is the one in joules or electronvolts provided that the appropriate value of the Planck constant h is also used.)

$$\lambda = (6.63 \times 10^{-34} \text{ J s}) \times (3.00 \times 10^8 \text{ m s}^{-1}) / 2.76 \times 10^{-17} \text{ J}$$

$$\lambda = 7.21 \times 10^{-9} \text{ m}$$

Since this result was obtained using the approximate relationship given in Equation 1.6, the answer needs to be expressed to one significant figure. Thus the typical wavelength of a photon that will be emitted by the corona is $7 \times 10^{-9} \text{ m}$ or 7 nm.

This example has shown that the thermal emission from the corona will occur at a wavelength which is about a hundred times less than visible wavelengths. Does electromagnetic radiation with such a tiny wavelength exist? If so, what is it? Once again, it's time for some more background science (Box 1.3).

BOX 1.3 THE ELECTROMAGNETIC SPECTRUM

In terms of the electromagnetic wave model, introduced in Box 1.1, visible light spans a range of wavelengths from approximately 400 nm to approximately 700 nm. Electromagnetic waves with wavelengths outside this range cannot, by definition, represent visible light of any colour. However, such waves do provide a useful model of many well-known phenomena that are more or less similar to light. For example, everyone is familiar with radio waves; we all rely on them to deliver radio and TV programmes. Radio waves are known to have wavelengths of about 3 cm or more; their well-established properties include the ability to be reflected by smooth metal surfaces and to travel through a vacuum at the same speed as light. Both of these properties, and many others that could have been

quoted, are also exhibited by electromagnetic waves of the same wavelengths. Thus we can say that radio waves can be 'well modelled' by electromagnetic waves with wavelengths greater than about 3 cm. Indeed, because the relatively long wavelength of radio waves makes their wave-like nature so obvious, it is really rather pedantic to talk about an electromagnetic wave *model* of radio waves at all. Most people would happily accept the statement that radio waves *are* electromagnetic waves.

The wide range of phenomena that can be modelled by electromagnetic waves is illustrated in Figure 1.36. As you can see, the full **electromagnetic spectrum**, as it is called, ranges from long wavelength **radio waves**,

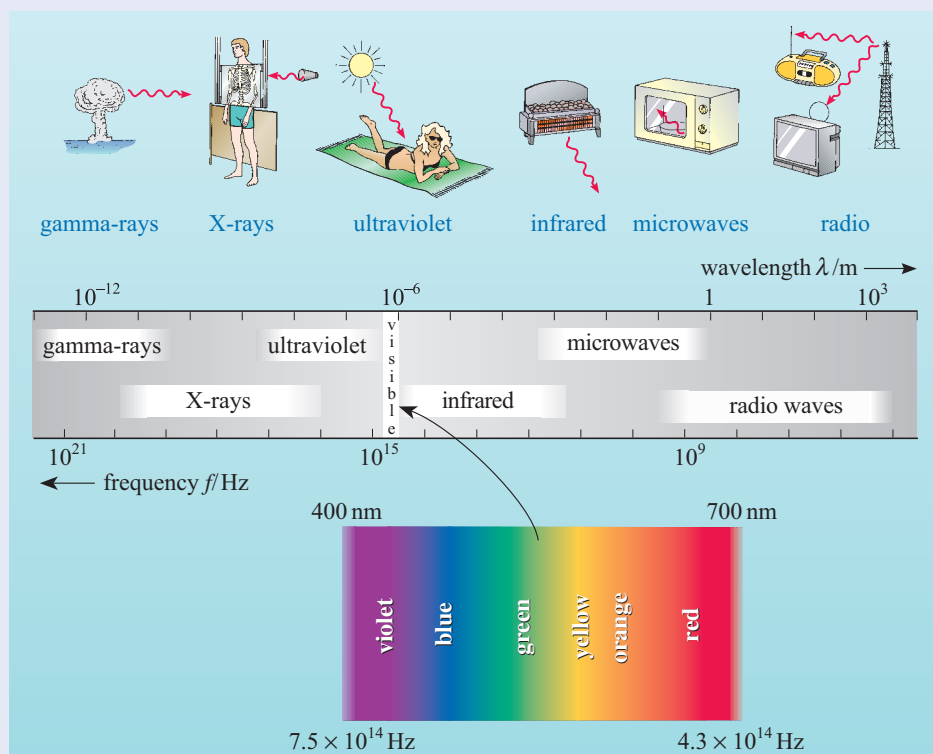


Figure 1.36 The electromagnetic spectrum. Note that the frequency and wavelength scales are logarithmic. Note also that the ultraviolet (meaning 'beyond the violet') adjoins the visible violet, and the infrared (meaning 'below the red') adjoins the visible red.

through **microwaves** and **infrared radiation**, across the various colours of **visible light** and on to such short wavelength phenomena as **ultraviolet radiation**, **X-rays** and **γ -rays**; (pronounced ‘gamma rays’). These various kinds of **electromagnetic radiation** arise in a wide range of contexts (as illustrated) but fundamentally they differ from one another only in the wavelength (or frequency) of the electromagnetic waves used to model them.

The boundaries of the various regions are deliberately vague; scientists and technologists often draw the divisions somewhat loosely.

The photon model of light that was introduced in Box 1.1 can also be applied throughout the electromagnetic spectrum. In fact, when it comes to γ -rays, their very short wavelengths make it quite difficult to demonstrate their wave-like properties, and it is much more conventional to speak of them as though they were particles.

Of course, the true situation is that all forms of electromagnetic radiation are, at the present time, most

accurately described by the quantum theory mentioned in Box 1.1. No part of the electromagnetic spectrum ‘really’ consists of waves or particles but any part may, under the appropriate conditions, exhibit wave-like or particle-like behaviour; both possibilities are encompassed by the quantum theory.

Many of the ideas in Section 1.3.2 concerning spectroscopy and sources of light are also applicable to the entire electromagnetic spectrum. Every kind of electromagnetic radiation may come from thermal or non-thermal sources, and the continuous spectrum produced by a black-body source always extends beyond the wavelength range of visible light. (This last point is already implicit in Figure 1.23, which shows the Planck curves stretching into the ultraviolet and the infrared.) Similarly, Wien’s displacement law applies throughout the electromagnetic spectrum. Also, spectral lines may arise at any wavelength provided transitions of the appropriate energy exist to cause them. Such transitions are not necessarily confined to atoms; they may involve nuclei, molecules or many other systems.

QUESTION 1.12

Complete Table 1.1.

Table 1.1 A partially completed table showing the properties of six different electromagnetic waves. For use with Question 1.12.

Wavelength, λ/m	3×10^{-14}	6×10^{-10}	5×10^{-3}	
Corresponding frequency, f/Hz	1×10^{22}	5×10^{17}		3×10^7
Corresponding photon energy, ε/J	7×10^{-12}		4×10^{-23}	2×10^{-26}
Corresponding photon energy, ε/eV	4×10^7	4	2×10^{-4}	1×10^{-7}
Temperature, T/K , of a black body that has a peak in its spectrum at this value of λ		5×10^6	3×10^2	3×10^{-4}
Corresponding part of the electromagnetic spectrum	γ -ray	X-ray		radio wave

1.4.2 The Sun's electromagnetic spectrum and the corona

Figure 1.37 shows the electromagnetic spectrum of the entire Sun across a range of wavelengths from 10^{-13} m to 10 m, or at least it attempts to do so. Such a spectrum, i.e. one that covers a wide range of wavelengths is often called a **broadband spectrum**. The central region of the graph in Figure 1.37, roughly from 10^{-7} m to 10^{-4} m, which accounts for most of the emitted power (note the logarithmic scales), is not difficult to interpret. Emission in this region is dominated by the photosphere and, apart from some absorption lines that do not really show up on this scale, the solar spectrum is well approximated by the Planck curve of a black-body source at 6000 K. Provided the Sun is in a quiet state with relatively little activity taking place, the emission of microwaves and short wavelength (i.e. less than 10 m) radio waves would be as expected from the Planck curve. However, at longer radio wavelengths the emissions come predominantly from the hotter regions higher in the Sun's atmosphere. One consequence of this is that the observed diameter of the Sun increases as the wavelength of observation increases, another is that the 6000 K Planck curve provides a progressively less satisfactory approximation to the observed spectrum. Moreover, if there is a significant amount of activity on the Sun, as is often the case, the related bursts of radio emission may

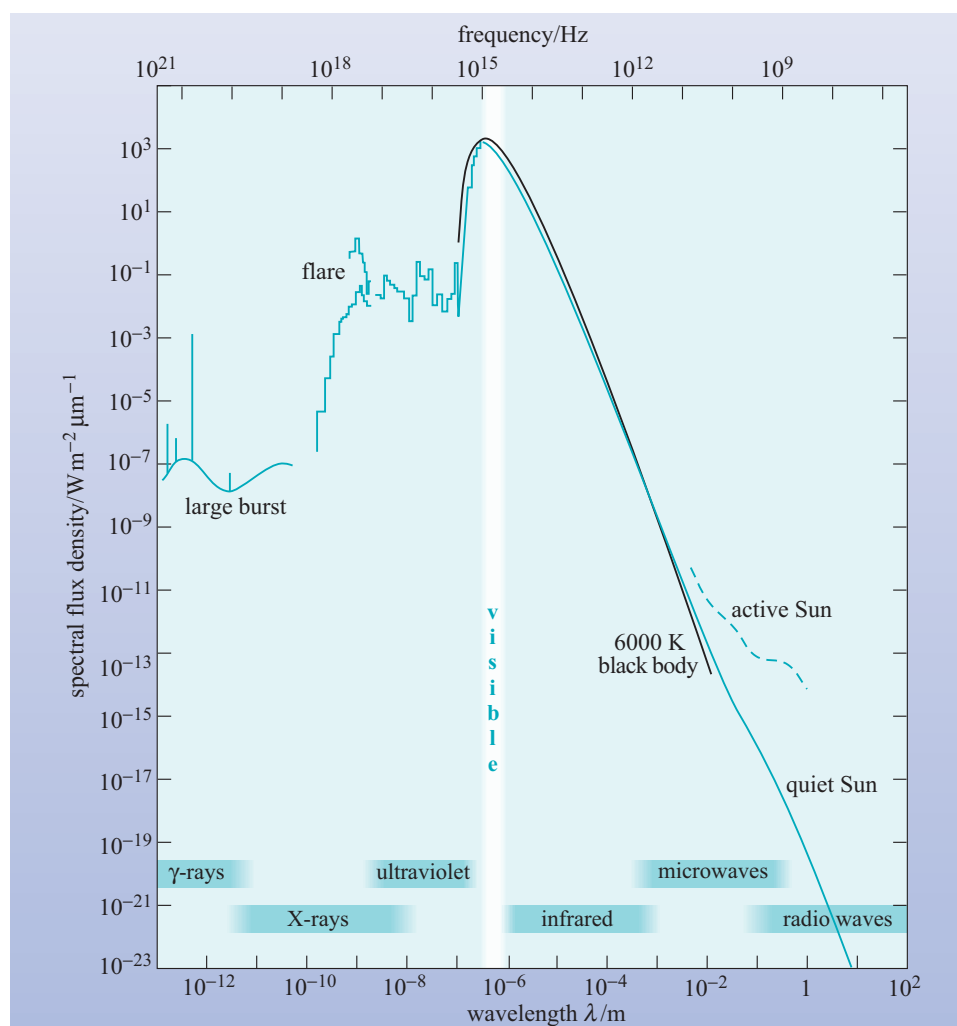


Figure 1.37 The solar electromagnetic spectrum. At any given wavelength λ , the height of the graph represents the amount of energy that would be received in 1 s by a detector, of area 1 m^2 , pointed directly towards the Sun, in a range of wavelengths of width $1 \mu\text{m}$ centred on λ , provided the detector was above the Earth's atmosphere. The shape of the graph is discussed in the text. The black line is the Planck curve of a black body at 6000 K. (Nicholson, 1982)

well dominate the spectrum over a broad range of wavelengths. This is especially true of explosive events called **solar flares**, one of the most energetic kinds of solar activity. We will return to discuss solar flares in more detail in Section 2.3.2. The existence of these outbursts makes the radio spectrum highly variable with time and difficult to represent on a simple graph.

Similar problems exist in the X-ray and γ -ray parts of the spectrum. The black-body emission from the photosphere falls off very rapidly with decreasing wavelength and is essentially negligible below a wavelength of about 10^{-8} m. Solar radiation is seen at shorter wavelengths, but it comes mainly from the hot coronal material and especially from the active regions where the temperatures are greatest. At the very shortest wavelengths (below 10^{-10} m) the individual X-ray and γ -ray photons each carry a relatively high energy, and it is quite normal to see spectra presented as ‘counts’ of the numbers of photons observed in various narrow wavelength (or energy) ranges in a given time.

Prior to the space age, it was impossible to obtain observational data about the broadband spectral behaviour of the Sun. The Earth’s atmosphere is a very effective absorber of certain wavelengths of electromagnetic radiation, with the consequence that Earth-based solar observations simply cannot be made for wide ranges of wavelength. Figure 1.38 provides a clear indication of the effectiveness of the Earth’s atmosphere as a barrier to solar radiation. The atmosphere is largely transparent to visible light and to certain longer wavelengths, particularly to narrow bands in the infrared and to radio waves, but everything else is effectively stopped. In order to overcome this barrier, many observations have to be made from high-altitude rockets and balloons, or satellites and space probes. Such ventures are complicated and costly, but they have been responsible for many of the most important recent advances in solar science (and in many other fields of astronomy).

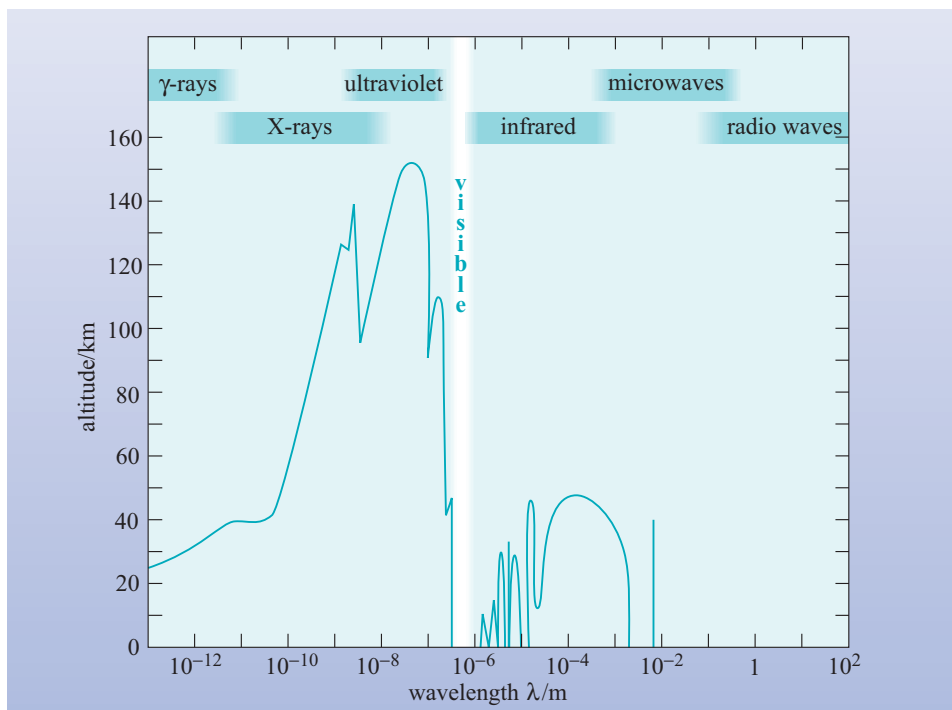
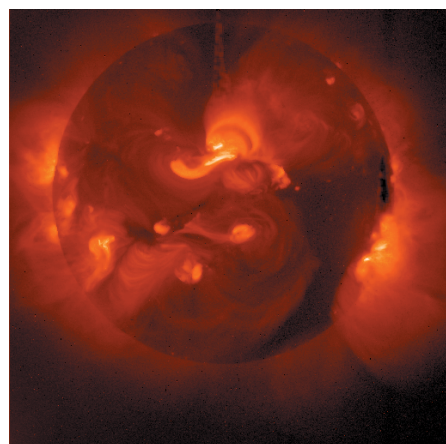
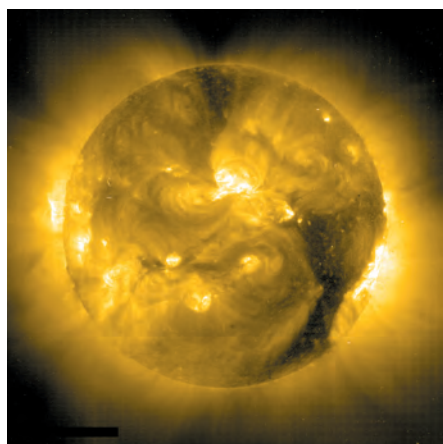


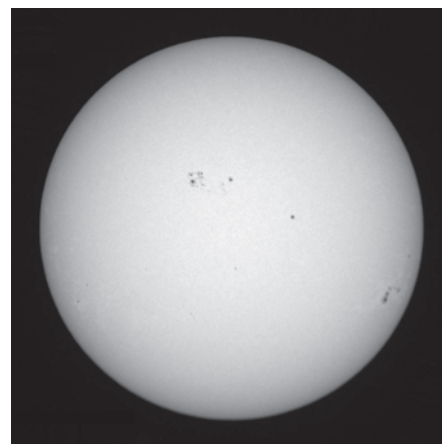
Figure 1.38 Effectiveness of the Earth’s atmosphere as a barrier to incoming solar radiation. For a given wavelength λ , the graph shows the altitude at which the spectral flux density (F_λ) due to the Sun is reduced to 50% of its value at the top of the atmosphere. (Nicholson, 1982)



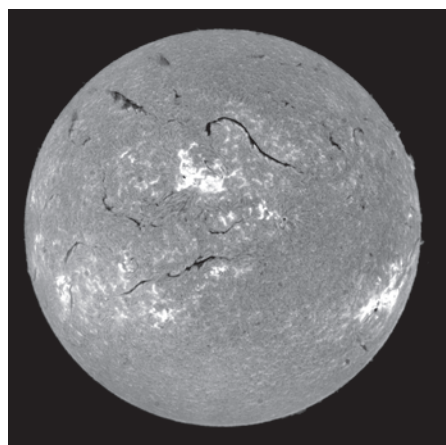
(a) X-ray (08:18 on 14 Dec 2001)



(b) Extreme ultraviolet (13:06 on 14 Dec 2001)



(c) White light (16:07 on 13 Dec 2001)

(d) $H\alpha$ (18:27 on 13 Dec 2001)

(e) Radio (23:05 on 13 Dec 2001)

Figure 1.39 The Sun at various wavelengths.

(a) X-ray ($\lambda = 0.3$ to 5 nm),
 (b) extreme ultraviolet ($\lambda = 28.4$ nm),
 (c) the whole visible spectrum,
 (d) $H\alpha$ ($\lambda = 656.3$ nm),
 (e) radio ($\lambda = 1.8$ cm). These images were all taken within a 24-hour period on 13 and 14 December 2001. ((a) Yohkoh (ISAS); (b) SOHO (ESA and NASA); (c), (d) Big Bear Solar Observatory; (e) Nobeyama Radioheliograph)

Five solar images, representing different parts of the electromagnetic spectrum, are shown in Figure 1.39. It should be clear from the above discussion of the solar spectrum that the images formed at very long and very short wavelengths will have been produced largely by radiation from the corona. The X-ray image (Figure 1.39a) is essentially confined to temperatures above 1.6×10^6 K and is particularly informative about coronal structure. Bright regions associated with solar activity are easy to see, but more intensive study reveals that the corona is generally composed of ‘loops’ of various sizes. These loops are magnetic in origin; they are known to result from the interplay of the highly ionized coronal gases and the Sun’s magnetic field. We will return to study these interactions in more detail in Chapter 2.

Another feature commonly seen in X-ray images is known as a **coronal hole**. Some idea of its nature can be obtained from Figure 1.40, which shows a sequence of X-ray images taken at intervals of 27 days from the orbiting Skylab space station in 1973. The rotation of the coronal hole – the dark boot-shaped object stretching from the pole to the equator – with the Sun is quite clear. The hole’s shape and size change with time, and sometimes it seems to fragment and then merge together again. Coronal holes are devoid of the large loops seen elsewhere in the corona: they are regions where the solar magnetic field opens outwards to interplanetary space rather than looping back on to the Sun. The existence of such regions is of great importance, since they are thought to be a major source of the **solar wind** –

a gusty stream of high-speed particles that spreads out from the Sun, carrying traces of the Sun's magnetic field with it. We will return to look in more detail at the origin and effects of the solar wind later in Section 2.4.

- Why were the Skylab images in Figure 1.40 taken 27 days apart?
- The Sun rotates; 27 days were required for the coronal hole to return to the centre of the field of view of the Skylab detector.

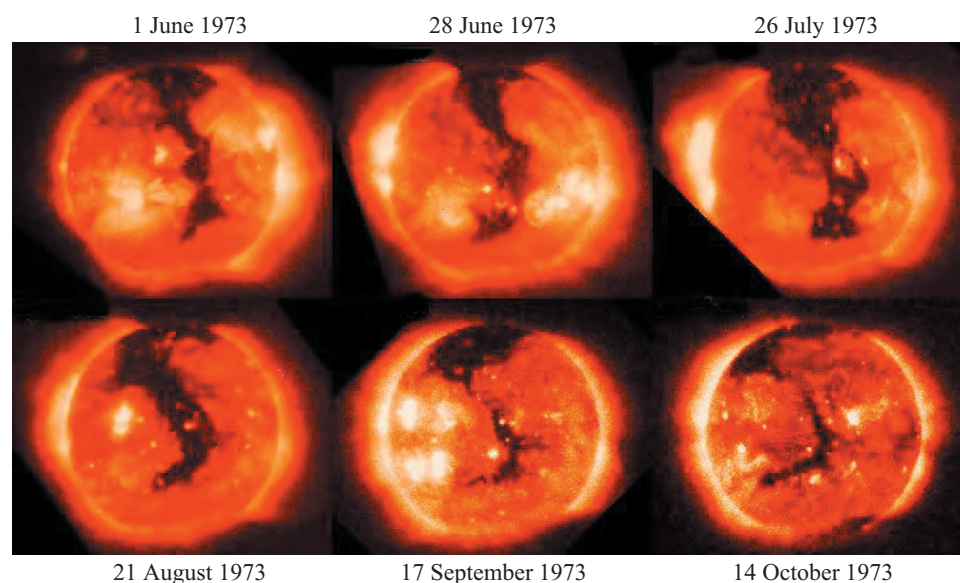


Figure 1.40 X-ray images of the corona taken at 27 day intervals (in 1973), and showing the evolution of a coronal hole. Such images must be obtained from space because the Earth's atmosphere absorbs electromagnetic radiation in this wavelength range. (Stanford University)

1.5 SUMMARY OF CHAPTER 1

Basic characteristics of the Sun

- The Sun is the closest star to the Earth.
- The solar luminosity (3.84×10^{26} W) is the rate at which energy is radiated by the Sun.
- The Sun emits radiation across the entire electromagnetic spectrum, from radio waves to γ -rays. The central wavelength range (10^{-7} m to 10^{-4} m), which accounts for nearly all the radiated power, is approximately described by a 6000 K black-body curve, and is dominated by photospheric emissions. The extreme wavelengths are more strongly related to solar activity and are often dominated by coronal emissions.

The nature of light

- Quantum theory provides the best available account of the nature of light, but simple models such as the electromagnetic wave model

$$c = f\lambda \quad (1.2)$$

and the photon model

$$\varepsilon = hf \quad (1.3)$$

are still of great value.

- A black-body source of radiation emits a characteristic continuous spectrum. The wavelength at which such a curve attains its maximum is determined solely by the temperature of the source, in accordance with Wien's displacement law:

$$(\lambda_{\text{peak}}/\text{m}) = \frac{2.90 \times 10^{-3}}{(T/\text{K})} \quad (1.4)$$

- For *any* thermal source of electromagnetic radiation (not necessarily a black-body source), the typical photon energy can be estimated by

$$\varepsilon \sim kT \quad (1.6)$$

- The absorption and emission lines found in the spectra of many sources of radiation are caused by electrons making transitions between states of different energies within atoms.
- The electromagnetic spectrum encompasses such seemingly diverse phenomena as radio waves, microwaves, infrared radiation, visible light, ultraviolet radiation, X-rays and γ -rays; all these forms of electromagnetic radiation can be modelled by electromagnetic waves of decreasing wavelength or by photons of increasing energy, respectively, under the appropriate circumstances.

The photosphere

- The photosphere is the visible 'surface' of the Sun. It is a thin semi-transparent shell of gaseous material about 500 km thick, 1.4×10^6 km in diameter and characterized by a temperature of about 6000 K.
- The photosphere exhibits limb darkening and differential rotation. The rotation can be traced by sunspots, which also indicate the level of solar activity.
- The Sun is the only star sufficiently close to Earth to permit the detailed study of small-scale phenomena, such as sunspots and granulation.

The chromosphere

- The Sun's chromosphere is a patchy layer of gaseous material, a few thousand kilometres thick, that overlies the photosphere. About 75% of the mass of the chromosphere is made up of hydrogen; helium accounts for most of the remainder.
- Detailed studies of solar spectra provide information about the chemical composition of the chromosphere and its physical condition (temperature, number density, etc.). Such studies show that the temperature of the chromosphere increases rapidly with height, changing by about 20 000 K over 2000 km.
- Chromospheric spectra show a number of emission lines, notably an H α line at 656.3 nm. H α images of the Sun reveal details of the structure of the chromosphere and various signs of solar activity, such as plages and filaments.

The corona

- The outer atmosphere of the Sun is called the corona. It is tenuous and extensive, and separated from the chromosphere by the narrow, irregular transition region. The structure of the corona is highly variable and responds to the level of solar activity.
- Spectral signatures of highly ionized atoms reveal that temperatures of several million degrees are common in the corona. Accounting for these high temperatures is still a major challenge for solar scientists.

Questions

QUESTION 1.13

By reference to a calcium K image of the Sun, estimate the size of a plaque.

QUESTION 1.14

Would it be possible to identify the material that makes up a black-body source by examining its spectrum?

QUESTION 1.15

The energy levels of the hydrogen atom (labelled E_1 , E_2 , E_3 , etc. in Figure 1.26) have energies that are well approximated by the formula

Note that all the energies are given negative values; E_1 is the lowest.

$$(E_n/\text{eV}) = -\frac{13.6}{n^2}$$

where n is the energy subscript, 1, 2, 3, etc. (Note that the photon energy is given in terms of electronvolts, so take care in the following question to use the Planck constant in the appropriate units.)

- Insofar as you can, confirm that the absorption occurring at the wavelength of the $\text{H}\alpha$ line can be attributed to transitions from E_2 to E_3 .
- Which transition accounts for the $\text{H}\beta$ absorption line? (*Hint*: attempt this by trial and error, with n no larger than 5. A useful skill to develop is that of bracketing an answer by guesswork and then homing in on it.)
- Which transition would account for an emission line at the same wavelength as the $\text{H}\beta$ absorption line?
- Why is there no line corresponding to the transition from E_1 to E_2 in the spectrum of Figure 1.28?

QUESTION 1.16

In addition to the green emission line due to ionized iron, the corona produces two other prominent emission lines in the visible part of the spectrum: a yellow line at 569.4 nm due to calcium atoms that have lost 14 electrons, and a red line at 637.4 nm due to iron atoms that have lost 9 electrons.

- How can the ions responsible for the yellow and red lines be represented symbolically?
- Calculate the difference in energy between the relevant energy levels for each of the two ions responsible for the yellow and red lines. Express your answer in electronvolts.

QUESTION 1.17

A thermal source of radiation has a temperature of 5×10^6 K.

- Calculate the (approximate) wavelength at which this source would emit electromagnetic radiation.
- If this source is not only a thermal source, but is also a black-body source, calculate the wavelength at which the peak of emission occurs.

(c) Explain why the answers to parts (a) and (b) differ, and state whether or not the answers are consistent with one another.

QUESTION 1.18

At visible wavelengths, images of the Sun exhibit the phenomenon of limb darkening, as described earlier. At radio wavelengths the opposite effect, limb brightening, is seen. Explain the origin of limb brightening. (*Hint*: recall that temperatures in the solar atmosphere increase with height.)
